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Star-Planet Interactions

Evry Schatzman School 2019

Lionel Bigot, Jérôme Bouvier,
Yveline Lebreton, Andrea
Chiavassa, and Agnès Lèbre, Eds

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Preface

The exoplanet revolution that began nearly three decades ago with the discovery of the first planet around a Sun-like star, has drastically changed our understanding of the formation of extra-solar systems. Today, we face an extreme diversity and complexity of these systems that can only be understood through the intimate link that exists between the planets and their host stars like tides, winds, and magnetic interactions. We are now in a position to detect planets down to Earth analogues thanks to space-borne missions like Kepler, TESS, and soon PLATO, and ground-based instruments installed on large telescopes like ESO/HARPS, CAHA/CARMENES, and CFHT/SPIRou. The detection and characterization of telluric planets is one of the main challenges of modern astrophysics. The difficulty lies in both detecting them and characterizing their properties and atmospheric composition, while the planetary signal is contaminated, if not drowned into the "noise" of the stellar activity. In this respect, the advent of a new and exciting new generation of instruments, like the James Webb Space Telescope, and ground-based instruments like CRIRES+ at VLT and in a near future PCS or ANDES at the E-ELT, will open a new area in exoplanet study with the detection of possible life tracers.

The 2019 edition of the Evry Schatzman School (EES 2019) of the CNRS/INSU French National Program for Stellar Physics (PNPS) was held in Aussois (France) from September 23 to 27¹. The aim of the school was to provide an overview of the new challenges related to exoplanetary systems with lectures on the characterisation of stellar variability which generates some "noise" for exoplanet detection (N. Meunier, G. Bruno), on star and planet interactions and their interplay with tides (S. Mathis), with magnetism (A. Strugarek), and on the impact of stellar winds on their companions (C. Johnstone). The challenges brought by exoplanets require an in-depth understanding of these stellar processes that are beyond the standard models of stellar structure and atmospheres. Another quest for astronomers is the characterisation of the atmospheres of exoplanets and, hopefully, the detection of biomarkers. These two topics were covered during the school through lectures on the definition and conditions of habitability (M. Turbet), and on the state-of-the-art techniques of detection of atoms and molecules in the atmosphere of exoplanets by transmission spectroscopy (M. Brogi). The school additionally offered hands-on sessions on tools used to overcome the impact of stellar noise, like Gaussian Processes (S. Aigrain), and the code SOAP to handle brightness features such as spots and plages on the stellar disk to estimate their impact on planet detection (M. Oshagh). All the material (slides of the talks and codes) are freely available on the website of the school. All the lecture notes in this book are also available on Arxiv.

We warmly thank our colleagues for the quality of the lectures and discussions during EES 2019; the students for their active participation with posters, presentations, discussions, and overall their enthusiasm; our sponsors: Centre National de la Recherche Scientifique (CNRS), Programme National de Physique Stellaire (PNPS) and Planétologie (PNP), Université de la Côte d'Azur (UCA), ERC SPIDI (grant agreement 742095), the Lagrange Laboratory and the Observatoire de la Côte d'Azur (OCA); the local organizing committee: C. Delobelle, S. Rousset, and I. Lapassat for managing the administrative and coordination aspects. We are also very much grateful to the staff of the Paul Langevin center for their warm welcome and all the efforts they put to make our stay fruitful and pleasant during that week, despite the on-going renovation of the center. We were glad to bring back the EES school at its original place in Aussois, where Evry Schatzman and Annie Baglin started this series², 30 years before the EES 2019.

L. Bigot, J. Bouvier, Y. Lebreton, A. Chiavassa, & A. Lèbre, the Editors

¹<https://ees2019.sciencesconf.org>

²<https://www.pnps.cnrs.fr/ees.html>



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Scientific and organising committees

Scientific committee

- L. Bigot (Chair)** - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France
J. Bouvier (co-Chair) - Institut de Planétologie de Grenoble, Grenoble, France
Y. Lebreton - LESIA, Observatoire de Paris, PSL Research University, CNRS, Sorbonne Université, Univ. Paris Diderot, Sorbonne Paris Cité, Meudon 92195, France and Univ Rennes, CNRS, IPR (Institut de Physique de Rennes)-UMR 6251, F-35000 Rennes, France
A. Chiavassa - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France
A. Lèbre - Laboratoire univers et particules de Montpellier, Montpellier, France

Local committee

- L. Bigot (Chair)** - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France
A. Chiavassa - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France
C. Delobelle - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France
I. Bailet - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France
S. Rousset - Université Côte d'Azur, OCA, CNRS, Lagrange, F-06304 Nice, France

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List of participants

1. **Jérémy Auir**, CEA - Département d'Astrophysique / Laboratoire AIM
2. **Suzanne Aigrain (Lecturer)**, Department of Physics, University of Oxford
3. **Lionel Bigot (Chair)**, Lagrange, Observatoire de la Côte d'Azur
4. **Jerome Bouvier (co-Chair)**, Université Grenoble Alpes, CNRS, IPAG, 38000 Grenoble, France
5. **Matteo Brogi (Lecturer)**, Department of Physics, University of Warwick
6. **Giovanni Bruno (Lecturer)**, Catania Astrophysical Observatory, INAF
7. **Tianqi Cang**, IRAP, Observatoire Midi-Pyrénées
8. **Andrea Chiavassa (SOC)**, Lagrange, Observatoire de la Côte d'Azur
9. **Martin Farnir**, STAR Institute, Université de Liège
10. **Eric Fossat**, Lagrange, Observatoire de la Côte d'Azur
11. **Guillaume Gaisné**, Institut de Planétologie et d'Astrophysique de Grenoble
12. **Lionel Garcia**, Université de Liège
13. **Kieran Hirsh**, CRAL, Univ-Lyon 1
14. **Colin Johnstone (Lecturer)**, University of Vienna - Dep. of Astrophysic
15. **Anand Utsav Kapoor**
16. **Laurel Kaye**, Department of Physics, University of Oxford
17. **Agnes Lebre (SOC)**, LUPM, Université de Montpellier
18. **Yveline Lebreton (SOC)**, LESIA, Observatoire de Meudon
19. **Louis Manchon**, IAS, Université Paris Sud
20. **Stéphane Mathis (Lecturer)**, CEA - Département d'Astrophysique / Laboratoire AIM
21. **Nadège Meunier (Lecturer)**, Institut de Planétologie et d'Astrophysique de Grenoble
22. **Belinda Nicholson**, Department of Physics, University of Oxford
23. **Mahmoudreza Oshagh**, Inst. for Astrophysics, Georg-August-Univ. of Göttingen
24. **George Pantolmos**, Institut de Planétologie et d'Astrophysique de Grenoble
25. **Camilla Pezzotti**, University of Geneva, Observatory of Geneva
26. **Jordan Philidet**, LESIA, Observatoire de Paris
27. **Charly Pinçon**, Université de Liège
28. **Kim Pouilly**, Institut de Planétologie et d'Astrophysique de Grenoble
29. **Noemi Roggero**, Institut de Planétologie et d'Astrophysique de Grenoble

- 30. **Melaine Saillenfest**, IMCCE, Observatoire de Paris
- 31. **Antoine Strugarek (Lecturer)**, CEA - Département d'Astrophysique / Laboratoire AIM
- 32. **Sophia Sulis**, Space Research Institute, Austrian Academy of Sciences
- 33. **Martin Turbet (Lecturer)**, University of Geneva, Observatory of Geneva, Versoix, Switzerland
- 34. **Arnaud Vericel**, Centre de Recherche Astrophysique de Lyon, Université Claude Bernard Lyon 1



Participants of the EES2019

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Physics of star-planet magnetic interactions

by A. Strugarek



Artist's impression of a hot Jupiter interacting with its host star through their magnetospheres. Copyright: NASA, ESA, and A. Schaller. Licensed under the Creative Commons Attribution-Share Alike 4.0 International. Source: https://commons.wikimedia.org/wiki/File:Artist%27s_impression_of_an_ultra-short-period_planet.jpg

Physics of star-planet magnetic interactions

Antoine Strugarek^{1,***}

¹AIM, CEA, CNRS, Université Paris-Saclay, Université Paris Diderot, Sorbonne Paris Cité, F-91191 Gif-sur-Yvette, France

Abstract. Magnetic interactions between a planet and its environment are known to lead to aurorae and shocks in the solar system. The large number of close-in exoplanets that have been discovered so far triggered a renewed interest in understanding magnetic interactions in other star-planet systems. Multiple magnetic effects were then unveiled, such as planet inflation or heating, planet migration, planetary material escape, and even some modifications of the host star apparent activity. Our goal here is to lay out the basic physical principles underlying star-planet magnetic interactions. We first briefly review the hot exoplanets' population as we know it. We then move to a general description of star-planet magnetic interactions, and finally focus on the fundamental concept of Alfvén wings and its implication for exosystems.

1 Introduction

1.1 The hot exoplanets population

In the long trailing wake of the discovery of 51 Peg b [26], more than 4000 exoplanets have been detected today. Space-based observatories such as **Kepler** and **Corot** have revolutionized this field by bringing a long-awaited wealth of exosystems, including unexpected massive planets on very short-period orbit (the so-called *hot Jupiters*). Complementing transit and radial velocity detection techniques, ground-based coronagraphy further allowed the direct imaging of distant exoplanets and revealed the existence of planets on very wide orbits (for a review, see [35]). The heterogeneous population of detected exoplanets is shown in Fig 1 as a function of the hosting star mass (vertical axis) and the orbital semi-major axis (horizontal axis), in units of stellar radius. The top panel presents the aggregated histogram of the known exoplanets as a function of their semi-major axis, which is heavily biased towards planets on compact orbits with a semi-major axis typically smaller than 20 stellar radii¹. Such close-in planets are expected to strongly interact with their host star through a variety of physical processes: they receive a strong irradiating stellar flux that affect their upper and lower atmospheres [18], they are subject to intense tides [24], and they orbit in a strongly magnetized medium. We will focus here on the latter and refer the reader to the cited reviews for the other types of star-planet interaction.

***e-mail: antoine.strugarek@cea.fr

¹Note that in the solar system, the semi-major axis of Mercury is about 83 solar radii

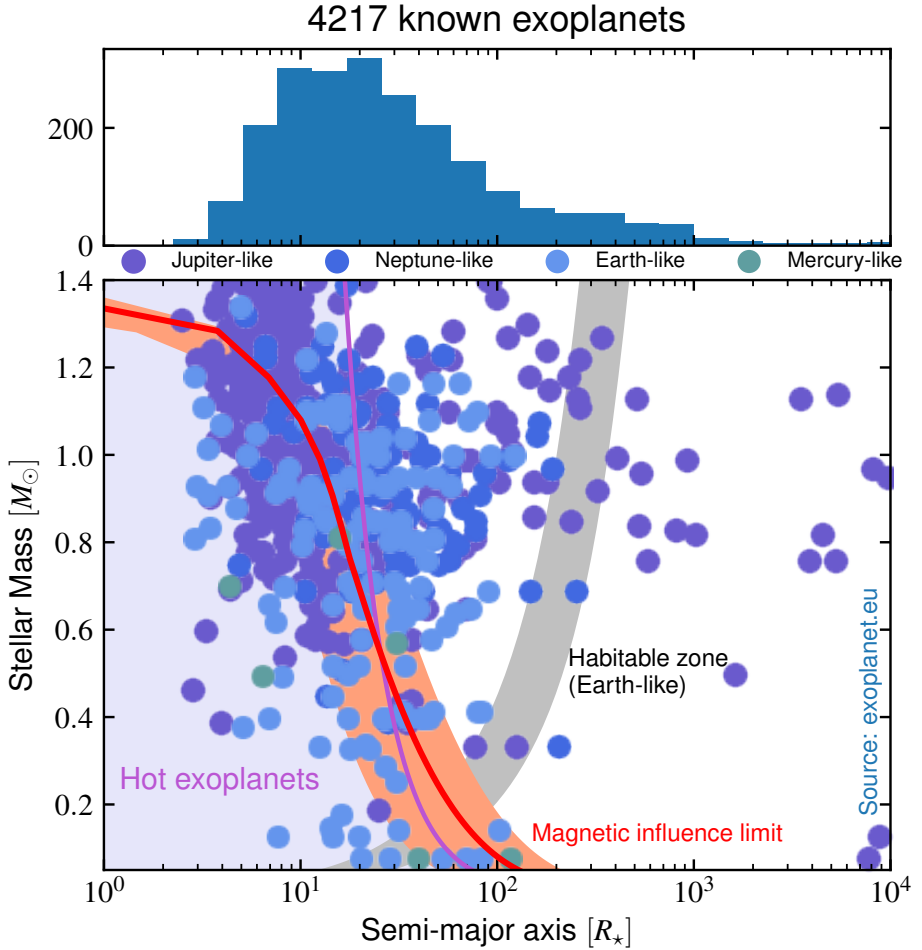


Figure 1: Exosystem distribution from exoplanet.eu. The bottom panel shows the distribution as a function host-star mass (vertical axis) and semi-major axis (horizontal axis), colored by the planet mass ($M_p \sin i$). The top panel shows the same distribution only as a function of the semi-major axis normalized to the stellar radius. The classical habitable zone [22] is shown as a grey band for an Earth-like planet. The $P_{\text{orb}} = 10$ days is labelled by the magenta line. Finally, a 'magnetic influence limit' line is shown in red (see § 2.1).

1.2 Observational status of star-planet magnetic interactions

The search for star-planet magnetic interactions has been a continuous history of positive and negative detections. The first firm detection of a signal that could be related to star-planet magnetic interactions was arguably reported by [38], following a seminal theoretical argument for the existence of star-planet magnetic interaction by [14]. They monitored at that time 5 promising star-planet systems, and detected a modulation associated with the hot Jupiter orbital period in the K band of Ca II of HD 179949. The other systems did not show any direct evidence of star-planet magnetic interactions at that time. Such signals have been observed since for different systems (see the review by [40]) such as famous compact

exosystems like 55 Cnc or HD 189733 [11]. The star-planet interactions tracers found in the stellar activity indicators have always been subject to a large variability with very clear signals at some epochs, and no detection at others (see for instance [10, 39]). We believe today that at least part of this hide & seek stems from the intrinsic variable nature of star-planet magnetic interactions. Star-planet magnetic interactions are controlled by the large-scale magnetism of stars, that can be topologically complex ([15, 37]) and time-variable due to eruptive events (e.g. coronal mass ejection, on a monthly time-scale) or magnetic cycles (on the yearly to decadal time-scale). As a result, star-planet magnetic interaction likely varies in nature and strength along planetary orbits as well as longer timescales.

Historically, tracers of star-planet magnetic interaction have been searched for in radio. Indeed, compact exosystems look alike planet-satellite systems in our own neighbourhood. The interaction of Io in the magnetosphere of Jupiter is known to lead to cyclotron-maser radio emissions at the poles of Jupiter [54]. At the same location, strikingly clear UV spots are present at the pole of Jupiter [13], tracing the footprints of the magnetic interaction of Jupiter with Ganymede, Io, and Europa. Until recently, no clear radio detection of star-planet magnetic interactions have been successfully obtained (for a recent review, see [55]). In 2019-2020, two new observations hinted toward these long-awaited radio emissions. The first one was a cautious suspicion of a radio emission in the τ Boo system [50]. The second one was the detection of coherent radio emission from the red dwarf GJ 1151 [51]. The latter is an indirect detection in the sense that the observed stellar radio emission is peculiarly intense compared to what one would expect from such a slowly-rotating star. The authors showed that the observed emission could be compatible with star-planet magnetic interaction with an Earth-like planet on a 1 to 5 days orbit, but we have not detected this planet so far. A firm detection of this hypothetical planet through transits or radial velocity would help strengthen our confidence in this detection scenario today.

2 Magnetic interactions in compact exosystems: overview

2.1 Sub- vs Super-alfvénic interaction

Planets in close-in orbits are generally thought to have their orbit circularized by efficient tidal interactions [41]. We will therefore consider here only planets in circular orbit, for the sake of simplicity. The characterization of star-planet interactions requires to estimate the plasma conditions along the orbit, that are set by the stellar wind. We will not detail stellar wind theory here and refer the reader to the previous Evry Schatzmann school proceedings by [8] and Johnstone (this volume) where the basics of magnetized stellar winds are introduced (the interest read can start by the seminal papers of [33, 52]). Let us summarize it quickly as follows. Cool stars like the Sun possess an external turbulent convective envelope that is the seat of an efficient dynamo process. The associated large-scale magnetic field then sculpts the environment of the star. In addition, turbulent photospheric motions are thought to excite magnetic perturbations, likely in the form of Alfvén waves, that propagate along this magnetic field and ultimately deposit energy upper the stellar atmosphere. In the case of the Sun, this leads to a very hot corona reaching more than a million Kelvin. This extremely hot corona then leads to the existence of a stellar wind, accelerating outwards and reaching super-sonic and super-alfvénic speeds. At the Earth orbit, the solar wind speed averages between 400 and 700 km/s, and reaches an alfvénic Mach number M_a between 9 and 10! In the context of exosystems, one can then wonder what is the speed of the stellar wind around close-in planets.

We can get a first estimate by computing a simple Weber-Davis [52] wind solution under the influence of rotation. Several alfvénic Mach numbers can be computed, each characteriz-

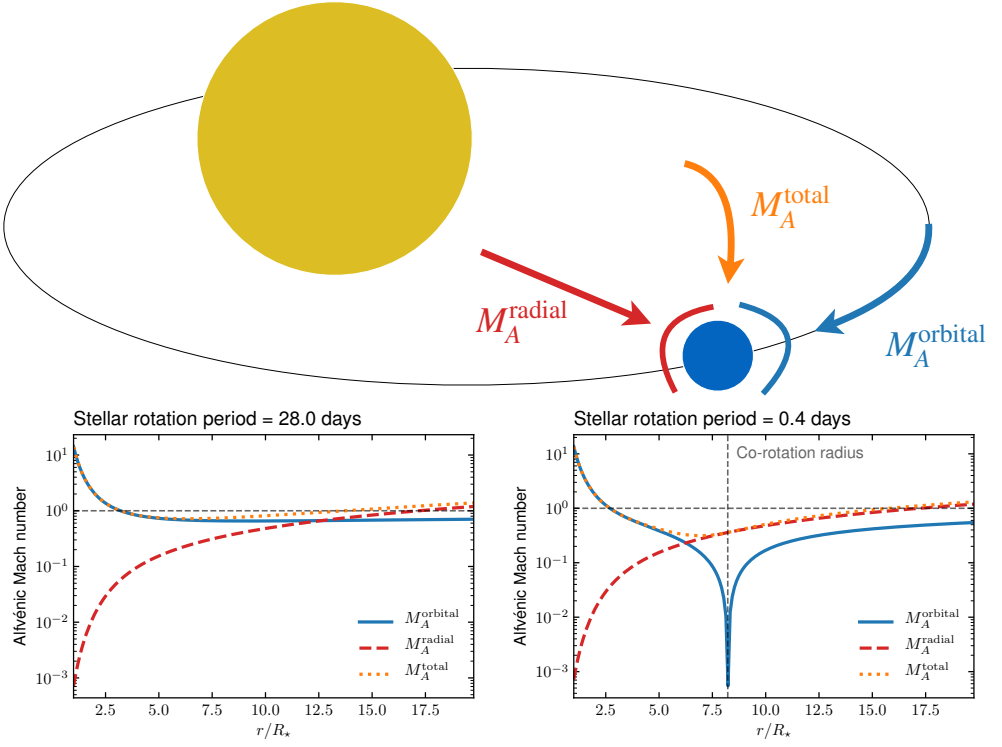


Figure 2: Schematic of alfvénic Mach numbers in compact exosystems. The top panel illustrates the star (yellow), the planet (blue) and its orbit (black), and the direction associated with the alfvénic Mach numbers (Eq. 1-3). The bottom panels show the values of these numbers as a function of orbital distance. The left panel corresponds to a solar-like star, and the right panel to a young, fast-rotating Sun.

ing different aspects of star-planet magnetic interactions. They are summarized in Fig. 2 and defined as follows

$$M_A^{\text{radial}} = \frac{v_r^w}{v_A}, \quad (1)$$

$$M_A^{\text{orbital}} = \frac{|v_\phi^w - v_{\text{kep}}|}{v_A}, \quad (2)$$

$$M_A^{\text{total}} = \frac{|\mathbf{v}^w - \mathbf{v}_{\text{kep}}|}{v_A}, \quad (3)$$

where $\mathbf{v}^w$ is the stellar wind velocity field, which we consider to be primarily in the radial star-planet direction (*i.e.* the accelerating component, denoted v_r^w) and in the orbital direction (due to stellar rotation, denoted v_ϕ^w). The Kepler velocity $\mathbf{v}_{\text{kep}}$ is assumed here to be purely in the azimuthal direction (circular orbit). Finally, $v_A = B_w / \sqrt{\mu_0 \rho_w}$ is the local Alfvén speed at the planet orbit based on the plasma properties in the stellar wind (where μ_0 is the vacuum magnetic permeability).

In the bottom panels of Fig. 2 we illustrate the different alfvénic Mach numbers for a solar-like star on the left and a fast rotating Sun (e.g. a young Sun) on the right. The classical alfvénic Mach number, based on the radial velocity of the stellar wind (Eq. 1), is shown by the red dashed lines. For the Sun, this simplified 1D model predicts an average alfvénic point ($M_A^{\text{radial}} = 1$) around 17 solar radii (i.e. 0.08 AU). The orbital Alfvénic Mach number is quite different as it is based on the relative orbital motion of the planet in the stellar environment. At first order, the corona of a star can be approximated to co-rotate with the star below the alfvénic point, and to rotate with a profile $v_\varphi^w \propto r^{-2}$ conserving angular momentum in the super-alfvénic region outside.

The orbital alfvénic Mach number M_A^{orbital} is shown in blue in both panels. It is super-alfvénic only very close to the star, due to a very fast keplerian motion. We see that it remains nevertheless sub-alfvénic for most of the domain for a solar-like star. In the case of an extremely fast-rotating star (right panel), the co-rotation radius (vertical dashed line) can be in the sub-alfvénic domain of the stellar wind. At the co-rotation radius, $M_A^{\text{orbital}} = 0$ by definition, which means that the magnetic interaction vanishes. Note that it is nevertheless a non-stable equilibrium of the star-planet system (see for instance [46]).

One can conclude from this crude analysis that many of the planets orbiting below $20 R_\star$ are likely to be in a sub-alfvénic interaction regime, with all Alfvénic Mach numbers below 1. We illustrate this by applying our simple 1D Weber-Davis wind model to stars with varying mass and rotation rate (typically here between 2 and 28 days). We obtain an estimate the Alfvénic transition, which appears as the red area along with the exoplanet distribution in Fig. 1. Note that some very close-in planets could still be in a super-alfvénic interaction state in their orbital direction, which still allows them to be magnetically connected to their host star in the radial direction. We now turn to the basic description of different magnetic connection scenarios in the sub-alfvénic regime.

2.2 Different regimes of sub-alfvénic interactions

A planet on a short-period orbit can be thought of as a perturber placed in a likely non-axisymmetric interplanetary medium. The perturbations it triggers can be decomposed on magneto-hydrodynamic (MHD) waves traveling away from the planet location (we will give the full description of this in § 3.1). A net Poynting flux is therefore channeled away from the planet by waves, along what is often referred to as Alfvén wings (see § 3). These alfvénic perturbations can travel back and forth between the star and the planet due to reflection along the magnetic field gradients. In the context of exosystems, a simple estimate of the ratio between the alfvénic travel time and the advective time-scale of the flow across the obstacle can be found in [49]. In essence, what is found is that for obstacles constituted by a planet and its magnetosphere, alfvénic perturbations generally do not have time to perform a back and forth travel between the planet and the low atmosphere of the star during the advective crossing time-scale². This situation corresponds to the so-called *pure Alfvén wing case* [31]. Other scenarios can be achieved, depending on the propagation time of these waves and on the conductive properties of the planet. For instance, [23] showed that planets with no magnetosphere and high internal conductivities could lead to the opposite situation where Alfvénic perturbations have time to perform multiple back-and-forth travels. The later case is generally referred to as the *unipolar inductor* interaction regime. A full description of the different interactions regimes can be found in [49], we refer the interested reader to this review and the references therein. While Alfvén wings have been observed on Earth (e.g. [12]) and

²Note that in general, the reconnection timescale is short enough to not play major role here in setting the type of star-planet magnetic interaction. It can nevertheless play a role in setting the absolute amplitude of these interactions.

for natural moons of solar system's giants planets (e.g. [7, 13], and references therein), the distinction between the different scenarios has proved to be difficult observationally. In all cases, though, two Alfvén wings always form in the $(\mathbf{B}_w, \mathbf{v}^w - \mathbf{v}^{\text{kep}})$ plane. Depending on the Alfvénic Mach numbers (§ 2.1), both wings can connect onto the host star; only one may do so while the other extends away toward the interplanetary medium, or both may head away from the host star. As a result, the knowledge of the magnetic configuration between the host star and the orbital path of the planet is mandatory to assess star-planet magnetic interactions.

3 Magnetic interactions in compact exosystems: the Alfvén wings

We now turn to the theoretical concept behind our understanding of star-planet interactions: the concept of *Alfvén wings* [30].

3.1 Pre-requisites on magneto-hydrodynamic waves

Let us recall here the properties of large-scale waves in magnetized plasmas. Neglecting rotation and stratification, three MHD modes can be identified: a slow MHD mode, a fast MHD mode, and pure Alfvén mode. The associated waves will be dubbed Alfvén wave (AW), slow MHD wave (SW), fast MHD wave (FW). Note that the term *magnetosonic* is generally used for purely perpendicular waves, we thus prefer to keep here the generic *MHD* denomination for these modes and waves. The derivation of the dispersion relation of these modes can be found in most good MHD textbooks (e.g. [17] to cite only one). We give only here the result and let the reader dig into such textbooks for their derivation.

We will write $\mathbf{k}$ the propagation vector of the wave, and ω its frequency. We will assume that at the studied scales the magnetic field can be considered to be homogeneous and we write $\mathbf{B}_0 = B_0 \hat{\mathbf{b}}$. The angle between $\mathbf{k}$ and $\mathbf{B}_0$ will be written θ . Finally, the classical Alfvén and sound speeds are defined as

$$v_A = \frac{B_0}{\sqrt{\mu_0 \rho_0}}, \quad (4)$$

$$c_s = \sqrt{\frac{\gamma P_0}{\rho_0}}, \quad (5)$$

where ρ_0 the plasma density, P_0 the plasma pressure and γ the ratio of specific heats. The dispersion equation for a pure Alfvén wave is

$$\frac{\omega}{k} = v_A \cos \theta, \quad (6)$$

and the dispersions for fast (+) and slow (-) MHD waves are

$$\begin{aligned} \left(\frac{\omega}{k}\right)^2 &= \frac{1}{2} \left[(c_s^2 + v_A^2) \pm \sqrt{(c_s^2 + v_A^2)^2 - 4c_s^2 v_A^2 \cos^2 \theta} \right] \\ &= \frac{v_A^2}{2} \left[(\tilde{\beta} + 1) \pm \sqrt{(\tilde{\beta} + 1)^2 - 4 \cos^2 \theta} \right], \end{aligned} \quad (7)$$

where $\tilde{\beta}$ is related to the plasma β through

$$\tilde{\beta} = \frac{c_s^2}{v_A^2} = \frac{\gamma}{2} \frac{2\mu_0 P_0}{B_0^2} = \frac{\gamma}{2} \beta. \quad (8)$$

Friedrichs diagram: phase velocity

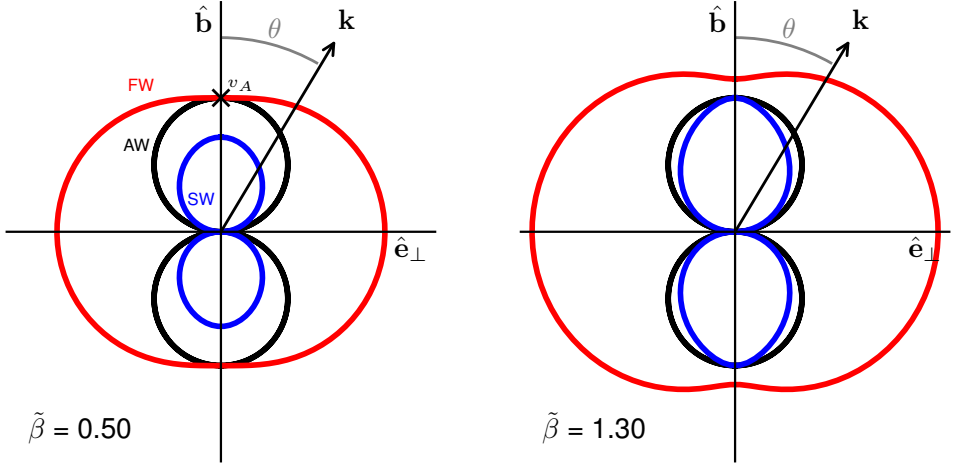


Figure 3: Friedrichs diagrams of the phase velocity of Alfvén waves (AW, black), fast MHD waves (FW, red) and slow MHD waves (SW, blue). The vertical axis is aligned with an ambient magnetic field, the horizontal axis represents any perpendicular axis. On the left panel the Alfvén speed is larger than the sound speed, on the right panel it is the opposite.

A graphical representation of Alfvén waves can help guide our intuition on their properties. In particular, Friedrichs diagrams can be built for these three fundamental waves. For a given frequency, we represent in Figure 3 the phase speed of the wave $\mathbf{v}_p = (\omega/k)\hat{\mathbf{k}}$ for all propagation directions $\hat{\mathbf{k}}$, where $\hat{\mathbf{k}}$ is the unit vector aligned with $\mathbf{k}$. In each panel, the magnetic field direction $\hat{\mathbf{b}}$ is represented by the vertical direction. The perpendicular direction $\hat{\mathbf{e}}_{\perp}$ is represented by the abscissa. In 3D, the Friedrichs diagrams are invariant by rotation along $\hat{\mathbf{b}}$. As a result, we represent these diagrams in 2D, bearing in mind that the full characteristics can be obtained by rotation along $\hat{\mathbf{b}}$.

We represent the phase speeds for the AW (black), the SW (blue) and FW (red). The AW and SW are degenerate for $\mathbf{k} = 0$ and fall back on what is sometimes called an *entropy wave*. Two illustrative cases are shown in Figure 3. On the left panel, $\tilde{\beta} < 1$ and the AW and FW have the same phase speed when $\mathbf{k} \parallel \hat{\mathbf{b}}$. On the right panel, $\tilde{\beta} > 1$ and the AW and SW have the same phase speed when $\mathbf{k} \parallel \hat{\mathbf{b}}$. We note that as $\tilde{\beta}$ increases, the sound speed overcomes the Alfvén speed and the AW and SW converge essentially to the same the phase speed portrait as the dispersion relation of SWs converges to the one of AWs.

The same diagram can be build for the group velocity of the waves. The group velocity is defined as

$$\mathbf{v}_g = \frac{\partial \omega}{\partial \mathbf{k}} = \hat{\mathbf{k}} v_p + k \frac{\partial v_p}{\partial \mathbf{k}}. \quad (9)$$

The first term corresponds to the parallel component of the group velocity (which is equal to the phase speed), and the second to perpendicular component of the group velocity. In the case of AWs, we directly obtain that only the parallel component remains and obtain

$$\mathbf{v}_g = v_A \hat{\mathbf{b}}. \quad (10)$$

Friedrichs diagram: group velocity

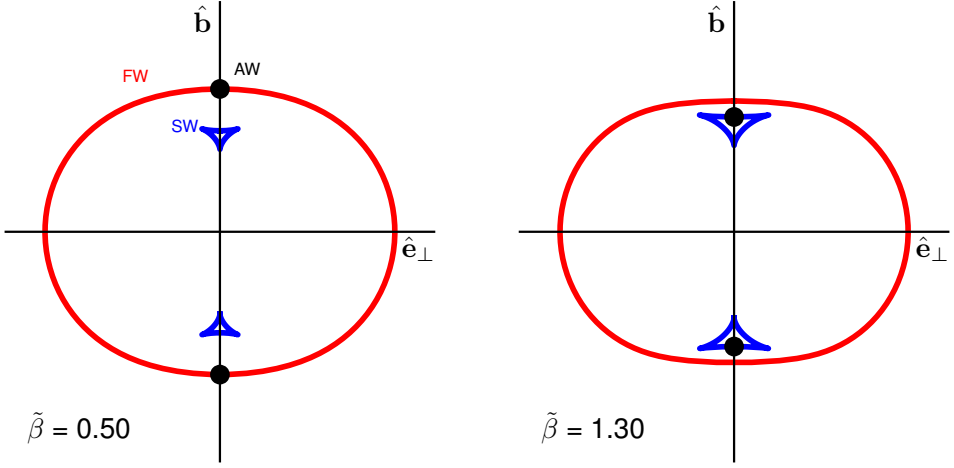


Figure 4: Friedrichs diagram for the group velocity of large-scale MHD waves. The layout is the same as in Fig. 3. Note that the diagram for Alfvén waves consists in the two black dots due to the degeneracy of their group velocity.

We immediately note an important result: the group velocity of pure AWs is degenerate as it does not depend on the wave vector $\mathbf{k}$. As a result any wave packet of pure AWs has to propagate along the magnetic field $\hat{\mathbf{b}}$. The same calculation can be followed for SWs and FWs, and one obtains

$$\mathbf{v}_g^{\pm} = v_p^{\pm} \left[\hat{\mathbf{k}} \pm \frac{\sigma \cos \theta \sin \theta}{2 \sqrt{1 - \sigma \cos^2 \theta} \left[1 \pm \sqrt{1 - \sigma \cos^2 \theta} \right]} \hat{\mathbf{n}} \right], \quad (11)$$

where we have introduced $\sigma = 4\tilde{\beta}/(\tilde{\beta} + 1)^2$, and $\hat{\mathbf{n}}$ is the direct unit vector perpendicular to $\hat{\mathbf{k}}$ in the $(\hat{\mathbf{b}}, \hat{\mathbf{k}})$ plane.

The group velocity indicates the propagation direction and speed of wave packets. It thus constitutes an important information to characterize how energy is channelled by wave ensembles. They are represented for AWs, SWs and FWs in Figure 4 with the same layout as in Figure 3. FWs are completely unfocused as their group velocity can take any direction. SWs have a more focused group velocity around $\hat{\mathbf{b}}$. This focusing is nevertheless less and less tight as $\tilde{\beta}$ increases. Finally, AWs are completely focused (black dots in Figure 4), as we already noted earlier. This will have major consequences on the existence of Alfvén wings, which we now turn to.

3.2 The concept of Alfvén wings

We will now apply the basic MHD concepts of Section 3.1 to the case of star-planet magnetic interactions. We want to characterize the stationary waves that can exist due to the existence of an obstacle in a magnetized flow. In this simplified view, the obstacle can be a conducting planet, or the ensemble of a planet with its magnetosphere. The detailed shape of the obstacle itself does not matter for the moment.

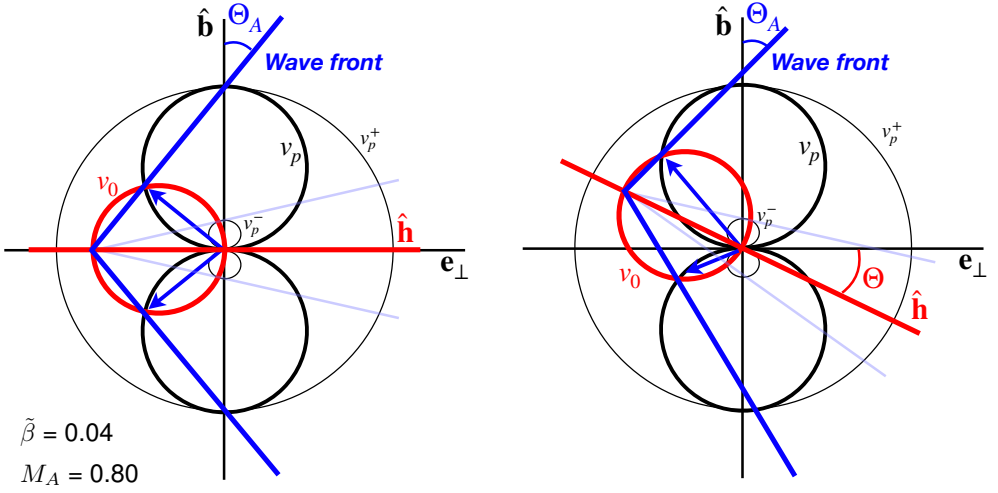


Figure 5: Friedrichs diagrams of the phase velocity involving a Doppler-shift associated with a large-scale flow impacting the obstacle. The Doppler-shift is labelled by the red circle. The direction of the incoming flow $\hat{\mathbf{h}}$ is given by the red line. The two panels show the same interaction case with a different impact angle Θ of the flow.

We will consider the frame at rest with the obstacle. Let us define $\mathbf{v}_0$ the flow velocity in this frame, such that the obstacle is subject to a stream $\mathbf{v}_s = v_0 \hat{\mathbf{h}}$ that deviates from being perpendicular to $\hat{\mathbf{b}}$ by an angle Θ . In this frame, a stationary wave solution requires that

$$\omega = \mathbf{k} \cdot \mathbf{v}_0, \quad (12)$$

that is naturally the Doppler shift in this frame. If one applies the same reasoning for Doppler shift that the one for Alfvén waves developed in previous section, we remark that such waves are described in the Friedrichs diagram by a circle of radius $v_0/2$ centered on $(-\cos \Theta v_0/2, -\sin \Theta v_0/2)$ in the $(\hat{\mathbf{b}}, \mathbf{e}_\perp)$ plane (again, in 3D this is actually a sphere by rotation along $\hat{\mathbf{h}}$). We can intersect this circle with the three Alfvén waves characteristics in the phase velocity Friedrichs diagram as shown in Figure 5. The possible stationary waves then depend on the most important control parameter for star-planet magnetic interactions: the (total) Alfvénic Mach number, defined here as

$$M_A = \frac{v_0}{v_A}. \quad (13)$$

On the left panel of Figure 5, we consider the canonical case $\Theta = 0$ for a sub-alfvénic interaction ($M_A = 0.8$). On the right panel, we illustrate the effect of the flow inclination angle Θ . The Doppler-shift is labelled by the red circle. We rightfully note that it actually intersects only the characteristics of the SWs and AWs. If the interaction was super-alfvénic ($M_A > 1$), the red circle would actually intersect the three characteristics. The Doppler, SW and AW circles all cross in the center. We find again the ‘entropy wave’, for which the wave vector $\mathbf{k}$ is perpendicular to $\mathbf{v}_0$ and the group velocity is equal to $\mathbf{v}_0$. These waves form a queue behind the obstacle, in the direction of the impacting flow.

We have seen in Section 3.1 that the group velocity of AWs is degenerate. As a result, thanks to the Friedrichs diagram analysis we know that pure AW wave packets will propagate along the so-called Alfvén characteristics (blue lines in Figure 5) defined as

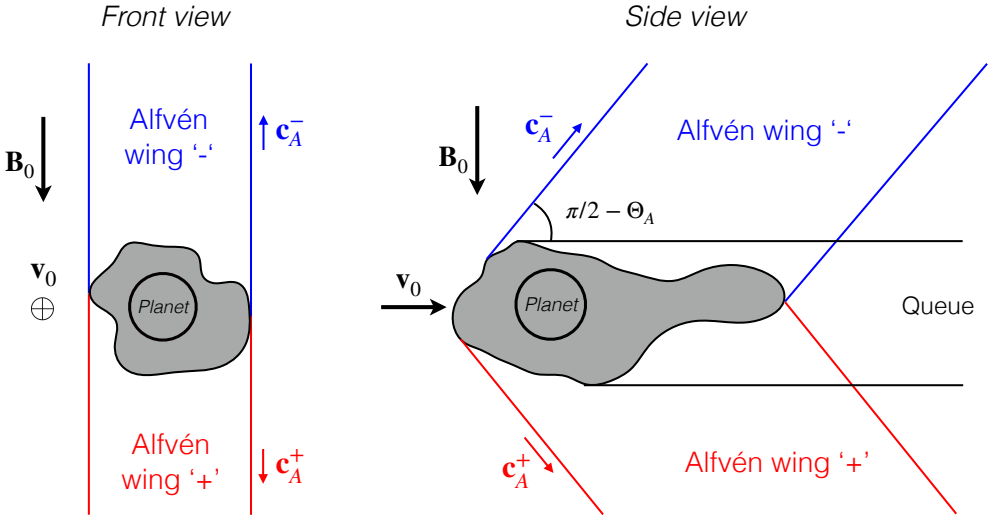


Figure 6: Schematic of the concept of Alfvén wings. The obstacle, illustrated by the blackened areas, can consist of a planet and its hypothetical magnetosphere. A front (left panel) and side (right panel) are shown. The two wings are labelled in red and blue.

$$\mathbf{c}_A^\pm = \mathbf{v}_0 \pm \mathbf{v}_A. \quad (14)$$

Finally, the case of SWs is interesting. Alike the pure AWs, they show an intersection with the Doppler shifted red circle in Figure 5. Nevertheless, their group velocity is non-degenerate. As a result, any excited SW wave packets will occupy a larger and larger volume as we move away from the obstacle. The energy and momentum transport by these waves will therefore be unfocused and are generally ignored in the context of star-planet or planet-satellite interactions.

We can sketch the characteristics of sub-alfvénic star-planet interaction deduced from this simple Friedrichs diagram analysis, as shown in Figure 6. The Alfvén wings always live in the $(\mathbf{b}, \mathbf{h})$ plane. The inclination angle between the Alfvén wings and the ambient magnetic field, Θ_A , can be geometrically derived from Figure 5 to obtain

$$\sin \Theta_A = \frac{M_A \cos \Theta}{\sqrt{1 + M_A^2 - 2M_A \sin \Theta}}. \quad (15)$$

To summarize, a first analytical description of Alfvén wings yields the following two robust conclusions:

- Only pure Alfvén waves and slow MHD waves are excited by the obstacle in the subsonic flow. Between the two, only pure Alfvén waves packets propagate in a focused manner, which leads us to consider only this wave population in the energetic discussion hereafter (§3.3)
- The overall shape of the Alfvén wings does not depend on the specifics of the obstacle. It is fully determined by the flow impacting the obstacle and the ambient magnetic field in the flow.

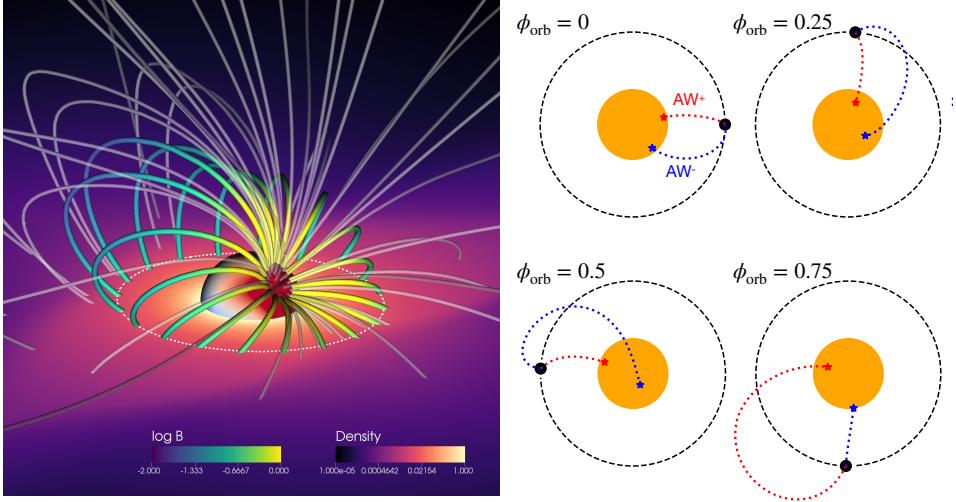


Figure 7: Alfvén wings in the Kepler-78 system [45]. The left panel is an illustrative view of the corona of Kepler-78 deduced from a 3D, MHD modelling using the observed magnetic map of Kepler-78 ([29]). The right panels show the projection of the two Alfvén wings on the ecliptic plane for four different orbital phases ϕ_{orb} . Figures are adapted from [45] with permission of the AAS.

The latter has a particularly interesting consequence: one can predict the shape of Alfvén wings solely based on the structure of a stellar corona and the orbital motion of an exoplanet. We show the application of the Alfvén wings theory to the Kepler-78 exosystem in Figure 7 (for more details see [45]). On the left, a polytropic wind model is used to derive the plasma properties of the corona of Kepler-78, based on observed Zeeman-Doppler-Imaging magnetic map observed by [29]. The orbit of Kepler-78b is shown by the dashed white circle. On the right, we apply the Alfvén wings theory to predict how wings change along the orbit of the planet. We show the projected trace of north (in red) and south (in blue) wings on the ecliptic plane. We observe that as the planet orbit the AW shape and sub-stellar points (red and blue star symbols) dramatically change due to changes in magnetic connectivity in the stellar corona.

3.3 Energetics of Alfvén wings

Now that we have understood how to characterize the shape of Alfvén wings, let us turn to their amplitude. How much energy is involved in magnetized star-planet interactions? The energy flowing through the wings is characterized by the Poynting flux $\mathbf{S}$ defined as

$$\mathbf{S} = \frac{\mathbf{E} \times \mathbf{B}}{\mu_0}. \quad (16)$$

Assuming ideal magneto-hydrodynamics, the electric field is simply $\mathbf{E} = -\mathbf{v} \times \mathbf{B}$. The projection of $\mathbf{S}$ along the Alfvén characteristics in the plasma reference frame is therefore

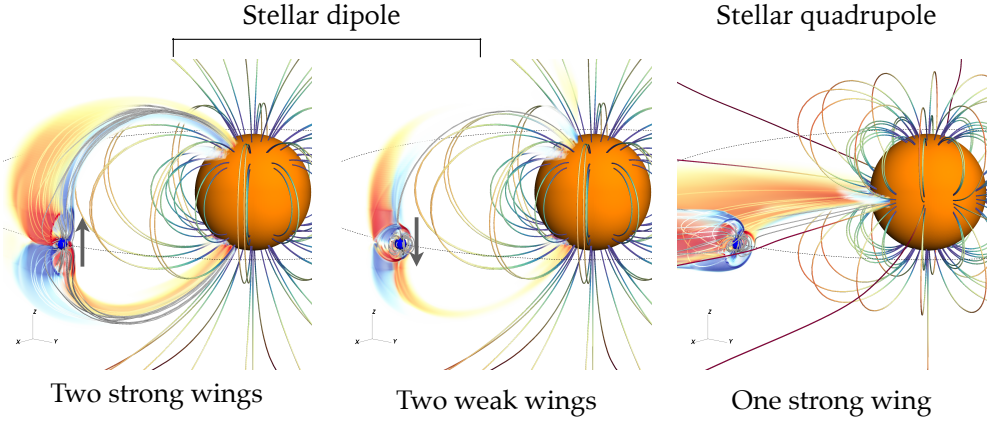


Figure 8: Illustration of the effect of the magnetic topology on Alfvén wings, from the 3D numerical simulations of [47]. The two left cases are the same except for the orientation of the planetary field (grey arrow), showing a modulation of the wings in **amplitude** (the red/blue volume renderings show the current system of the Alfvén wings). The right panel is the same except for the stellar magnetic field that is quadrupolar, showing a change in the **shape** of Alfvén wings. Figures are adapted from [47] with permission of the AAS.

$$S_{AW} = \mathbf{S} \cdot \frac{\mathbf{c}_A^\pm}{|\mathbf{c}_A^\pm|} = \frac{v_0 B_0^2}{\mu_0} \sin \Theta_A \cos \Theta. \quad (17)$$

Equation 17 gives a robust, maximum estimate of the power involved in the magnetic interaction. It corresponds to the case where the perturbations propagating in the wings are sufficiently large so that the incoming flow is completely halted inside the wings.

In general, the exact Poynting flux depends on two important aspects of star-planet magnetic interaction: (i) the details of the interaction inside the grey area in Fig. 6: what are the plasma, magnetic, and conductive properties of the obstacle? and (ii) the size of the effective obstacle. The latter determines the thickness of the Alfvén wings, over which one has to sum to estimate the total power involved in the interaction. We illustrate these aspects in Fig. 8 where we show 3D simulations for three different magnetic topologies [47]. The Alfvén wings are highlighted by the volume rendering of the currents in blue/red (negative/positive). All three cases are equivalent: in the two first cases only the orientation of the planetary field is reversed, while in the third case the stellar magnetic field is changed from a dipole to a quadrupole of similar amplitude. We strikingly see in these renderings that the relative topology of the magnetic field has an important influence on the strength of the wings, while their overall shape is control by the stellar atmosphere.

[36] have developed an analytical model to calculate the total Poynting flux over Alfvén wings, assuming the obstacle is magnetized. They found, for simple geometries,

$$\mathcal{P} = \int_{AW} \mathbf{S} \cdot d\mathbf{A} \simeq 2\pi R^2 \bar{\alpha}^2 (1 + M_A^2 - 2M_A \sin \Theta)^{1/2} S_{AW}, \quad (18)$$

where R is the effective radius of the obstacle, and $\bar{\alpha}$ quantifies the conductive properties of the obstacle³. These results were later confirmed and extended to more complex geome-

³ $\bar{\alpha} = \Sigma_P / (\Sigma_P + \Sigma_A)$, where Σ_P is the Pedersen conductance in the planet's ionosphere and $\Sigma_A = (1/\mu_0 v_A)(1 + M_A^2 - 2M_A \sin \Theta)^{-1/2}$ is the Alfvén conductance.

tries through 3D numerical simulation [42, 48]. The maximal power is obtained for high conductance in planet’s atmospheres for which $\bar{\alpha} = 1$. In the solar system, $\bar{\alpha}$ varies from a few 10^{-2} to about 1 for the different moons of Jupiter and Saturn [21, 36]. The conductive properties of exoplanets are nevertheless largely unknown as of today. They depend for instance on the complex photo-chemical reactions in their upper atmosphere, and hence on both the irradiation received by the planet, and the planet atmospheric composition. Both are expected to change significantly along stellar evolution. Ab-initio calculations are today being performed to characterize (among other aspects) the conductive properties of exoplanets’ atmosphere (*e.g.* see the Kompot code, [20]).

Assuming a maximized interaction, the total Poynting flux $\mathcal{P}$ requires the knowledge of the plasma environment around a planet along its orbit. [36] considered a simple Parker-like wind and applied their formalism to the ensemble of known exoplanets at that time. The largest uncertainty lies here in the lack of knowledge of the key stellar properties (rotation period, magnetic field) in the sample. For this review I revisited the updated ensemble of **Kepler**-observed exosystems while using the wind-modelling strategy developed by [1]. This strategy can be applied for the stars with a detected rotational period ([28]: a stellar Rossby number can then be estimated based on the scaling law derived from 3D numerical simulation by [9]. The stellar magnetic field and the subsequent wind is then computed following the approach proposed by [1] that satisfies all the observational constraints on stellar wind available to date. The resulting estimated Poynting flux (Eq. 18) assuming a planetary surface field of 4G for these systems is shown in Fig. 9. In spite of the large scatter due to the large variety of star-planet systems considered here, we see a clear overall trend $\mathcal{P} \propto P_{\text{orb}}^{-7/2}$ for the compact exosystems in the **Kepler** field. We highlight two populations in Fig. 9: exosystems in a very likely sub-alfvénic interaction regime (red circles), and exoplanets in a likely sub-alfvénic interaction regime (green squares). The former population is found to be in a sub-alfvénic interaction within all the error-bars of the models and observations used here (see [1] for more details on these error-bars). The latter population (green squares) is predicted to be sub-alfvénic only for part of our estimated magnetic field and coronal properties of the central star. It is nonetheless instructive to observe that sub-alfvénic interactions are highly unlikely to occur for planets with an orbital period longer than five days.

The Kepler-70 system stands out in this diagram. It is composed of a hot sub-dwarf around which a super-Earth (Kepler-70b) orbit with an extremely short period. Our estimates predict a moderate stellar magnetic field, which translates into a quite weak star-planet magnetic interaction. As a result, small orbital periods do not necessarily imply strong star-planet magnetic interactions. This is also illustrated by the large vertical spread in Fig. 9: planets with the same orbital period can lead to absolute Poynting fluxes varying by more than 3 orders of magnitude (*e.g.* for $P_{\text{orb}} \sim 2$ days) depending on the spectral, rotational, and magnetic properties of their host star.

The case of Kepler-78 is also highlighted in the top left of Figure 9. The Poynting flux estimate was carried out using our empirical scaling law with no *a priori* knowledge of the stellar magnetic field. It nonetheless provides an estimate that is very close to the one obtained when carrying out fully 3D MHD simulations of Kepler-78, based on an observed magnetic maps with Zeeman-Doppler Imaging (see Fig. 7 and [45]).

We have illustrated the theoretical derivation of the energetics of star-planet magnetic interaction with the **Kepler** sub-sample of exoplanets. This sub-sample is particularly interesting because asterosismology allows us to have access with good precision to the fundamental stellar parameters required to properly assess magnetic interactions. Many other exosystems as for instance HD 1897333 or 55 Cnc [11] also present promising configurations for sub-alfvénic interactions. Dedicated, 3D models of such systems are today required to properly assess both the energetics and the ephemerids of star-planet interactions [45]. One can hope

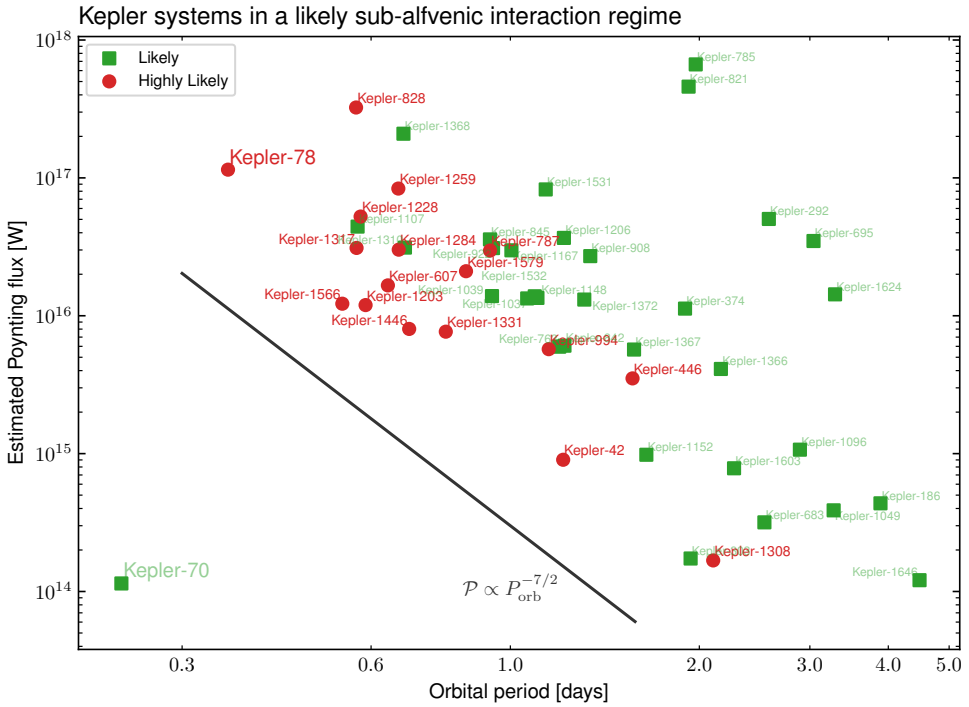


Figure 9: Estimated Poynting flux power (Eq. 18) for the Kepler sub-sample of [27] as a function of the orbital period of the planet. The red dots label the planets that are always predicted to exhibit star-planet magnetic interactions within all the error-bars of the model. The green squares label the planets that are predicted to exhibit such interactions only for the most optimistic error-bars.

that the coming years will see the definitive confirmation of star-planet magnetic interactions detection thanks to a combination of advanced 3D models supporting this interpretation of peculiar observational features correlated with exoplanets' orbital periods.

3.4 Angular momentum transfer and the secular evolution of stellar systems

The mere existence of Alfvén wings implies an energy transfer that also contains in part an *angular momentum transfer*. Indeed, an orbiting planet interacts with its environment and therefore is subject to an equivalent drag force. For the solar-system planets, the stellar wind is not quite dense enough for this force to matter at all in their orbital evolution. But it may not be systematically the case for close-in exoplanets orbiting in a dense and strongly magnetized stellar corone.

The magnetic torque felt by close-in exoplanets was systematically parametrized through numerical simulations [42, 48]. We show the example of one such simulation in the left panel of Fig. 10, for one particular magnetic topology (see [19] and [47] for more details on the effect of topology). A clear magnetic torque originating from the magnetic tension around the planet (green line) produces most of the total net torque (grey line) felt by the orbiting planet. The main unexpected finding of these studies was that this net magnetic torque could make close-in planets migrate on a time-scale of hundreds of millions years. Henceforth, magnetic

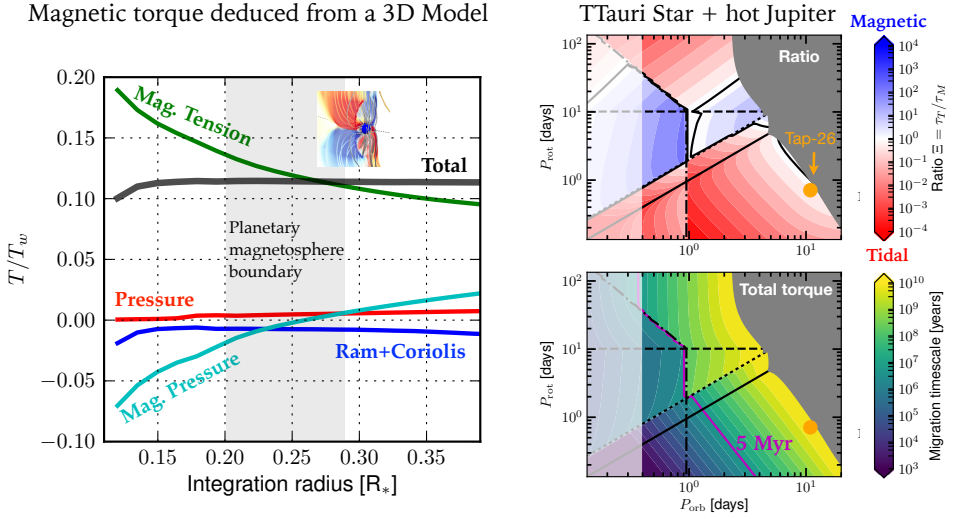


Figure 10: Illustration of magnetic torques due to star-planet magnetic interactions. The left panel shows an angular momentum budget around the orbiting exoplanet from 3D numerical simulations [47]. The main magnetic torque is dominated by the magnetic tension contribution in green. The right panel shows the comparison between tidal and magnetic migration torques in a $(P_{\text{rot}}, P_{\text{orb}})$ diagram for TT-Tauri – hot Jupiter exosystem (see [44] for more details). The bottom right panel shows the total (tidal+magnetic) migration timescale in a logarithmic color-scale for the same system. Figures are adapted from [47] and [43] with permission of the AAS.

torques could in principle play a role in shaping the population of close-in exoplanets. This situation typically arises mainly for strongly magnetized dwarf stars, or early in the evolution of a solar-like star. For main-sequence slow-rotating stars, magnetic torques play a more negligible role: for instance, the fastest migration timescale found in the **Kepler** sub-sample studied in Fig. 9 is predicted for Kepler-828 with a timescale of approximately 2 Gyrs.

Nevertheless, it is important to consider magnetic torques for ensemble evolution of compact exosystems, particularly during their early evolution. A systematic comparison between tidal and magnetic torques can be carried out thanks to the developments in both theoretical fields over the last decade. We have carried out such a comparison in [44] and showed that in general both physical phenomenons add up to make orbital migration more efficient. We found that magnetic torques were likely dominating the migration process for fully-convective stars and young stars, while tidal torques were likely dominating otherwise. We summarize this instantaneous comparison in the right panel of Fig. 10, for a typical TTauri star - hot Jupiter system (alike Tap-26, see [53]). We see that the torque can be dominated by either tidal torques (red regions in the top panel) or magnetic torques (blue regions) in a $P_{\text{rot}}-P_{\text{orb}}$ diagram. The bottom panel shows the migration timescale resulting from the sum of tidal and magnetic torques (see [44] for more details).

It is today mandatory to go beyond the instantaneous approach to understand how compact star-planet systems jointly evolve. Any orbital migration corresponds to an angular momentum exchange with the star, that ultimately affects the stellar rotation period, and thus its dynamo action and the subsequent evolution of the system. This is a challenging en-

deavour as it involves the complex stellar secular evolution along with non-linear physical phenomena such as stellar winds, and tidal and magnetic star-planet interactions. A seminal exploration was carried out simultaneously [5, 34, 56] exploring the angular momentum budget of a two-body star-planet system along stellar evolution, including tidal interactions. Recently, [16] followed a similar approach based on state-of-the-art stellar evolution tracks produced by the STAREVOL code [3], and advanced stellar wind [25] and tidal interactions [6, 32] modelling. In this study they focused on the hot exoplanets population predicted by such models for young open clusters such as Pleiades. A similar approach was also used with the ESPEM code [4] to predict which star-planet systems were the most likely to undergo a coupled orbital-rotational evolution phase. Star-planet magnetic torques yet have to be included in these theoretical explorations. This tool can ultimately be used to produce synthetic exoplanet populations, as we recently demonstrated [2].

Such studies are today highly valuable in the context of **Kepler** and **TESS** (and in the future **PLATO**) as the number of the observed exosystems keeps growing fast. They will allow to disentangle the relative importance of the different physical effects in shaping the population of exoplanets thanks to dedicated, careful cross-comparisons between observed and synthetic samples.

4 Conclusions

We have presented here the basics behind our physical understanding of star-planet magnetic interactions. Several *grey* areas nevertheless remain to be explored and understood. Among them, I would like to highlight in particular that we have almost not discussed the details of the obstacle itself, *i.e.* the planet and its hypothetical magnetosphere/ionosphere/etc... While the topology of star-planet magnetic interactions is largely controlled by the star itself, their strength is sensitive to the detailed properties of the obstacle. What are the conductivities in the different layers of the atmosphere and interior of exoplanets? This question cannot be trivially answered today, as exoplanets can have wildly different chemical reactions compared to solar system planets, that can receive a much different ionizing flux from their star, and they can be subject to much stronger tidal interactions with their host star and/or other planets in the system.

In spite of these hurdles that have to be overcome on the theoretical side, the characterization of star-planet magnetic interactions should allow us to probe the magnetism of exoplanets. This grand goal is worth such efforts, as today we can only characterize (with not so much precision) the magnetism of the few planetary bodies in the solar system. One may await a revolution in our understanding of planetary magnetism if the hunt for star-planet magnetic interactions succeed in the future. I hope this review will help motivate more and more young researchers to dive into the fantastic open world of star-planet interactions.

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Stellar variability in radial velocity

by N. Meunier

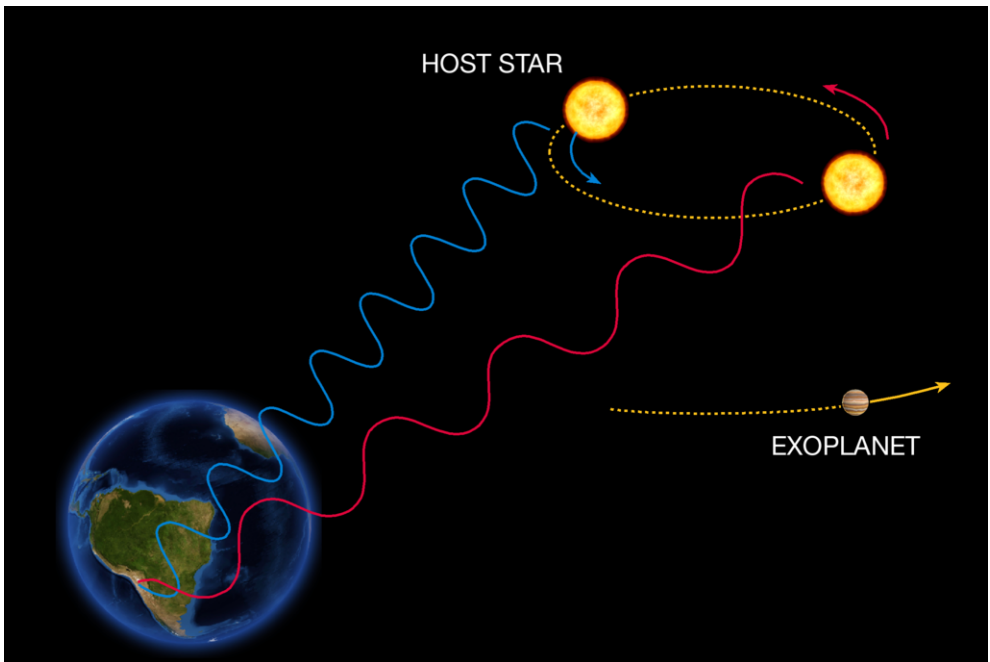


Illustration of the radial velocity method. Copyright: ESO. Licensed under the Creative Commons Attribution-Share Alike 4.0 International. Source: https://commons.wikimedia.org/wiki/File:ESO_wobble.png

Stellar variability in radial velocity

Nadège Meunier^{1,****}

¹Univ. Grenoble Alpes, CNRS, IPAG, F-38000 Grenoble, France

Abstract. Stellar activity due to different processes (magnetic activity, photospheric flows) affects the measurement of radial velocities (RV). Radial velocities have been widely used to detect exoplanets, although the stellar signal significantly impacts the detection and characterisation performance, especially for low mass planets. On the other hand, RV time series are also very rich in information on stellar processes. In this lecture, I review the context of RV observations, describe how radial velocities are measured, and the properties of typical observations. I present the challenges represented by stellar activity for exoplanet studies, and describe the processes at play. Finally, I review the approaches which have been developed, including observations and simulations, as well as solar and stellar comparisons.

1 Introduction

Stellar activity impacts the measurement of radial velocities (hereafter RV), which in turn affects the detectability of exoplanets using this indirect technique: this was recognized very early-on [148] after the detection of 51Peg b [99]. This lecture therefore focuses on integrated RV observed for stars other than the Sun, and on the effect of both magnetic activity (due to spots and plages) and variability due to photospheric dynamics at various spatial scales (from granulation to large scale flows), as well as the interaction between these two categories of processes. Asteroseismology as well as Doppler imaging are outside the scope of this lecture. We focus on F-M main sequence stars, with a bias toward old solar-type stars.

RV is an indirect technique in exoplanet studies, i.e. we observe the light coming from the star and not from the planet (Fig. 11): if the star impacts the RV measurement (by modifying the line shape or its position), then exoplanet detectability and characterization are also affected. This is true for several indirect techniques (RV, photometry, astrometry), which are complementary in terms of the orbital range they cover, but also in terms of the information they provide (for example, mass is obtained with radial velocities and radius with transit photometry, giving clues on the planet density). However, although RV are very sensitive to stellar activity, they are not the traditional way to study it: RV observations and surveys are usually biased toward the search for exoplanets (studies in which stellar activity is usually considered as “noise”), while several other methods are commonly used to study stellar activity such as chromospheric emission, photometry, or spectro-polarimetry.

Furthermore, most observables used to study stellar activity are strongly degenerate, as illustrated in Fig 12: for example, there is a strong degeneracy between spots and plages or

**** e-mail: nadege.meunier@univ-grenoble-alpes.fr

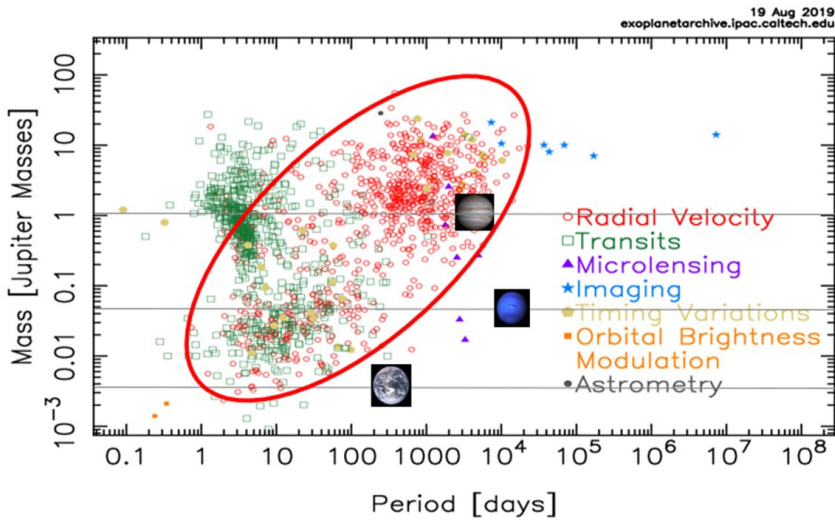


Figure 11: Planet mass vs. planet period (projected mass when observed with radial velocity) for different detection techniques. Jupiter, Neptune and the Earth are indicated for comparison.

between structure size and contrast when analyzing photometric light curves, while polarimetry is very sensitive to flux cancellation. RV measurements are no exception in that respect: they are sensitive to many processes, some of them subject to degeneracies. However, all these techniques are sensitive to different processes and are therefore complementary from the point of view of stellar activity. Combining different approaches and activity indicators is therefore very useful to obtain a complete view of stellar variability for such stars: RV in particular are sensitive to several processes not affecting other observables, which is a problem from the point of view of exoplanet detection, but also very interesting to study stellar processes related to magnetic activity or photospheric flows. Studying stellar activity using RV methods is also necessary to push the limits of exoplanet detections. Finally, a great wealth of knowledge has been obtained for the Sun: solar physicists are used to exploit the very high spatial resolution at our disposal, often without considering the integrated signal (except for irradiance and some helioseismology observations mostly), but leading to results of great interest to understand stellar RV. The content of this lecture therefore lies at the interface between stellar physics, solar physics and exoplanet studies.

The outline of the lecture is the following. I first introduce the context, how radial velocities are measured and affected, and what are the current challenges. Then I review the stellar processes impacting radial velocity measurements in detail and provide information about their typical amplitude and timescales. Finally, I describe the approaches which have been developed by different groups to deal with the challenge consisting in disentangling the planetary contribution from the stellar one: stellar and solar observations, correction techniques and simulations. I conclude with some perspectives for the next few years.

2 From the instruments to radial velocities

The RV technique needs stabilized spectrographs associated to high spectral resolution to be able to measure very precise RV on a long-term basis (> 10 years). Very stable instruments

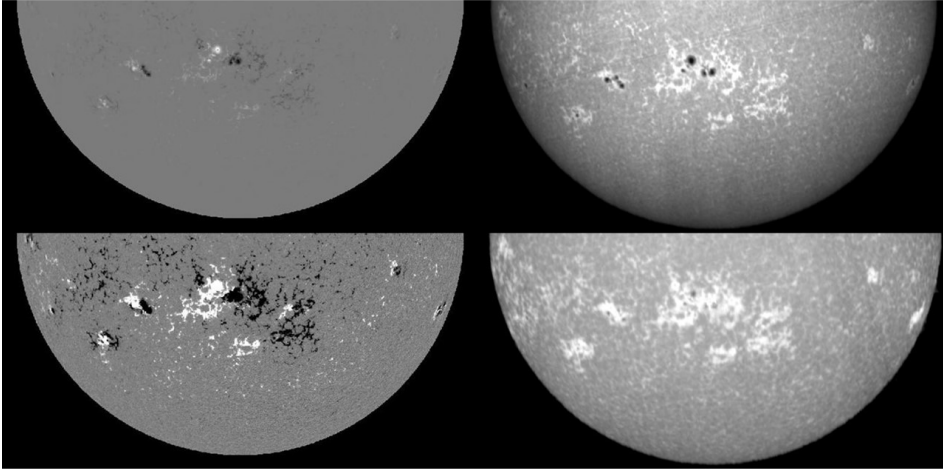


Figure 12: Solar magnetogram obtained by MDI/SOHO (left, with two different magnetic field saturation on the figure), and images of the solar photosphere (upper right panel) and chromosphere (lower right panel) from Meudon Observatory (BASS2000).

have been developed and implemented to be able to detect exoplanets, which are inducing a reflex motion of the star: this technique, based on the measurement of the position of spectral lines using various techniques, led to the first detection of an exoplanet around a main-sequence star in 1995 [99]. I present here a short overview of instruments and observations.

2.1 Instruments

Such spectrographs have been developed for the purpose of detecting exoplanets using RV techniques (this list is not exhaustive as many facilities exist) in two wavelength domains: first in the visible (e.g. Sophie at the 1.93m, Observatoire de Haute Provence, HARPS at the 3.6m in La Silla, HARPS-North at the TNG in La Palma, Espresso at the VLT in Paranal using one of the 8m UT or the 4 telescopes) and more recently covering the infrared domain as well (e.g. Carmones at the Calar Alto Observatory, Spirou at the CFHT, Hawaii, or the soon to come NIRPS in La Silla). Their main characteristics are: a high resolving power (in the 70k-190k), a very high long-term stability (with goals between 10cm/s and 1 m/s depending on the instrument and wavelength domain), and a good signal-to-noise ratio. They are characterized by their transmission, wavelength domain, and sampling. The S/N ratio on the spectrum depends on the spectral type and magnitude of the star, and on the order in the spectrum, but is typically in the 10-500 range. Reaching values down to 10 cm/s constitutes a huge instrumental challenge, as a displacement of the spectrum on a typical CCD detector represents a few atoms only!

2.2 From spectra to radial velocities

These instruments produce Echelle spectra, each of them typically recording the stellar spectrum as well as a reference spectrum used for the precise wavelength calibration (Th lines,

Fabry-Perot etalon, ...) for each order (for example 72 for HARPS). RVs are usually computed for each order separately and then combined. Different methods have been implemented to measure RV from these spectra. The most common one (for example used in the standard SOPHIE or HARPS pipeline) is based on a cross-correlation between the observed spectrum and a binary mask indicating the position of the expected lines for the spectral type of the observed star. The position of the maximum of this cross-correlation function (CCF), found by adjusting a symmetric Gaussian function, provides the stellar RV for this observation. Several groups have also developed complementary techniques, which are more suitable for certain types of stars, and in which the correlation function is computed with a reference spectrum instead of a binary mask, usually a median spectra derived from the whole time series of spectra for that star. This has been done in the Fourier space [30, 54] or in the temporal space [5, 6, 8], to adapt the algorithm to very massive stars with very few broad spectral lines and to M stars with many blended lines respectively. There are also prospects to compute the RV more globally using Gaussian Processes [138]. A few typical amplitudes give an idea of the difficulty of this task. Solar line widths are typically 2 km/s for example, and faster rotators have line widths higher than 100 km/s, while a HARPS pixel is about $0.01 \text{ \AA}$ (corresponding to 600 m/s at $5000 \text{ \AA}$). For comparison, detected planets have amplitudes of the order of the m/s up to a few 100 m/s for the most massive ones, while the Earth is below 10 cm/s only.

2.3 Observational strategies

Because RV surveys have been biased toward the search for exoplanets, the temporal sampling is not necessarily optimal for stellar activity. The sampling is usually very irregular, with a bad phase coverage of the stellar processes (for example of the rotational period). The sampling also affects the typical orbital periods which can be detected. This sampling results from various observational constraints such as telescope time allocation, the observability of the star, and the fact that few nights per year are available to observe a given star (visibility of the star and reasonable airmass to observe it when visible). Large surveys have been implemented, with data publicly available in some cases, for example on the ESO archive for HARPS: old FGK main sequence stars [98, 160], old main sequence M stars [16, 122], A-F stars [18, 19], young M-F stars [56, 78]. These surveys are however often biased toward the less active stars in the considered category, with the objective to minimise the effect of stellar activity and maximise exoplanet detection rates.

RV technique applied to exoplanet detection is not the only one to be affected by stellar activity, since this is also the case for photometric transits. However, although the stellar signal is present at all time in both cases, the planet contributes to the signal in a very different manner: it is always present in the case of RV, while this is the case only during the planetary transits for photometry, meaning that in the latter case, the light curves are most of the time free of the planetary signal, as shown in Fig. 13. It is therefore much easier to disentangle the planet signal from the stellar signal in this case (the transit depth can however be impacted by stellar activity, see lecture by G. Bruno, this volume). In addition, photometric light curves (from spatial missions such as CoRoT, Kepler, or TESS) are always very well sampled, while RV samplings are extremely irregular and sparse, as illustrated in Fig. 13 for two typical time series.

Stellar activity also affects the astrometric signal, but for massive planets to be detected with Gaia, the amplitude of the effect is completely negligible for solar-type stars. Future high-precision astrometric mission aiming to detect very low mass terrestrial planets will be sensitive to stellar activity however, but to a much lesser extent than radial velocities for similar stars and planets [79, 92, 113].

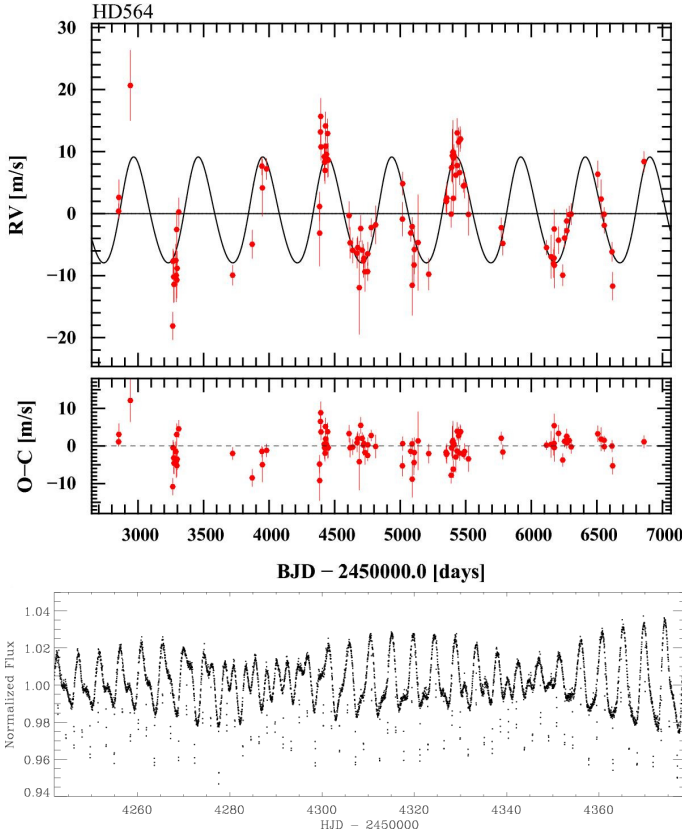


Figure 13: Upper panel: HARPS RV data for HD564, from [126] reproduced with permission ©ESO. Lower panel: example of a photometric time series with CoRoT, showing the stellar variability superimposed to the transit of CoRoT-2b, from [4] reproduced with permission ©ESO.

2.4 Complementary information: chromospheric emission

When searching for exoplanets, observables which are sensitive to stellar activity but not to the presence of a planet are routinely produced in order to help disentangling a possible planetary contribution from the stellar signal. They are computed from the same spectra used to obtain RV (and are therefore simultaneous): FWHM (full width at half maximum) and quantification of the line shape variability (for example the BIS, defined as the difference between the positions of the bisector at two different levels). Another very widely used indicator is a measure of the chromospheric emission, strongly related to magnetic activity, such as the $\log R'_{HK}$ indicator, representing the emission in the Ca II H and K lines (at 3933 and 3968 Å): a first indicator is defined by the emission integrated over the center of the line normalized by the continuum (S-index). The photospheric contribution is then subtracted, using a calibration depending on B-V in the 0.6-1.2 range [130] and more recently for M stars [7]. This calibrated flux is then corrected by the bolometric flux [130] to allow to compare stars of very different spectral types, which leads to the routinely produced $\log R'_{HK}$ indicator. The estimate of this indicator is likely to present uncertainties due to these calibrations, of the order of 0.05 for the average level of a given star [136], but its variability is more

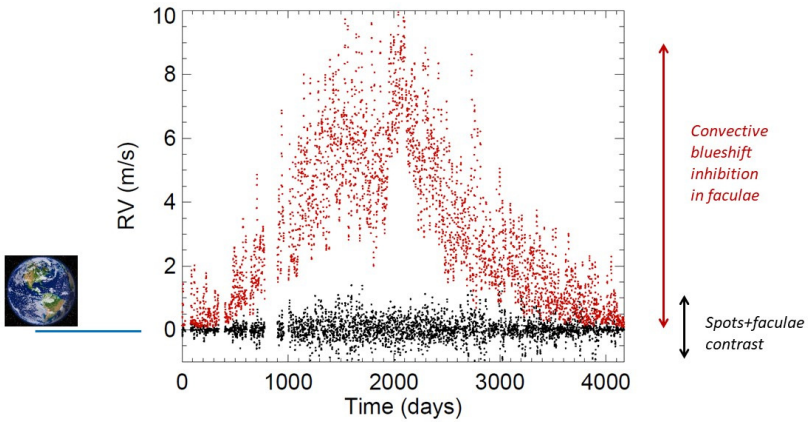


Figure 14: Reconstruction of the solar integrated RV due to magnetic regions, due to the spot and plage contrasts (black) and including the inhibition of the convective blueshift in plages (red) made by [105], compared with a Earth signal (~ 9 cm/s, smaller than the thickness of the blue line).

precise and very useful to identify the presence of magnetic activity. The chromospheric emission is strongly related to plages, i.e. bright magnetic area in the photosphere (and not to dark spots), as can be seen in solar images taken in the same wavelength band (BASS2000, <http://www.bass2000.obspm.fr/>), as shown in Fig 12 (lower right panel), or for example the high spatial observation around a small solar structure [57]. The chromospheric emission can also be derived from other chromospheric lines, for example $H\alpha$, but the correlation with the $\log R'_{HK}$ (as a function of time) is not always good [32], which could be due to stellar processes [104].

2.5 Effect of stellar variability on exoplanet studies in radial velocity: impact and challenges

Stellar variability, due to various processes (mostly magnetism and photospheric flows as well as processes related to both categories), varies at different timescales: they will be detailed in the next section. Their impact on exoplanet observations can take different forms:

- *It can look like an exoplanet*, and in the past led to exoplanet detection publications which were later invalidated. For example, the re-analysis [145] of the four planets detected around G1581 [97, 159] showed that planet d did not exist, and that the observed signal at its period was due to stellar activity. Another example is the publication of two Super Earths in resonance orbiting HD41248 using 62 points of public data [71], while the complete analysis by the PI of the observations [150] on many more points showed that one of the detections was due to stellar activity, while the other one did not exist. Most of recent publications take the presence of stellar activity carefully into account, using activity indicators as described in Sect. 2.4, although all mitigation techniques have limitations. This will be discussed in Sect. 4. The estimation of the false positive level is however a difficult issue [153].

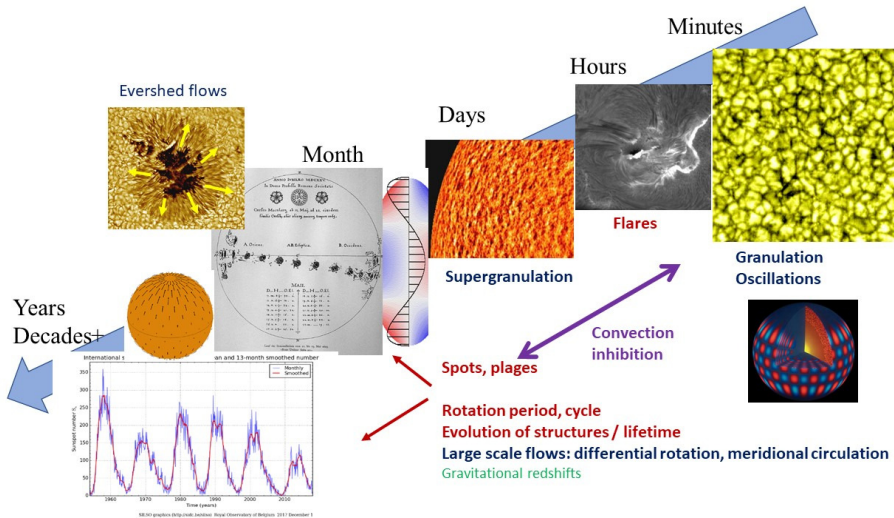


Figure 15: Overview of the different processes at play, due to magnetic regions (red), flows (blue), interaction between the two (purple), and gravitational redshift (green). Typical time scales correspond to the Sun. Granulation image: Pic du Midi Observatory; Flare: $H\alpha$ image, Big Bear Solar Observatory; Supergranulation: Dopplergram MDI/SOHO; oscillations: example of a mode from the GONG project, NSO; differential rotation: NASA/Marshall Solar Physics; spot drawings: Scheiner, 1625; Evershed flows: photosphere image from Vacuum Tower Telescope, NSO/NOAO; meridional circulation flows: Fig. 1 from [93], reproduced by permission of the AAS; spot number vs. time from SIDC.

- *It can hide an exoplanet* because of the strong additional signal. As an example, Fig. 14 shows the reconstruction of the solar RV signal during cycle 23 [105], leading to the conclusion that the inhibition of the convective blueshift in plages is dominating over the contribution due to the contrast of spots and plages: it has a long term amplitude of about 8 m/s, which is two orders of magnitude higher than the Earth signal, hence a huge challenge.
- *It can affect planet characterisation*, although this has been less studied. The estimation of the planet mass with RV techniques in transit follow-ups can therefore be noisier or biased. The presence of the stellar signal leads to significant uncertainties on the masses determined by RVs and possibly to some biases, which in turn will impact the estimation of the planet densities. It could also affect the Rossiter-McLaughlin effects, or planetary atmosphere characterization.

3 Stellar processes contributing to radial velocities

3.1 General overview

Stellar activity varies on different time scales. Figure 15 shows a summary of the different processes at play, where the indicated scales are indicative of the typical solar case: they may be quite different for other spectral types or young stars. These processes are due to magnetic structures, to flows (in particular in the photosphere), to interaction between magnetism and flows, and to gravitational redshifts. At the shortest timescales, oscillations and granulation

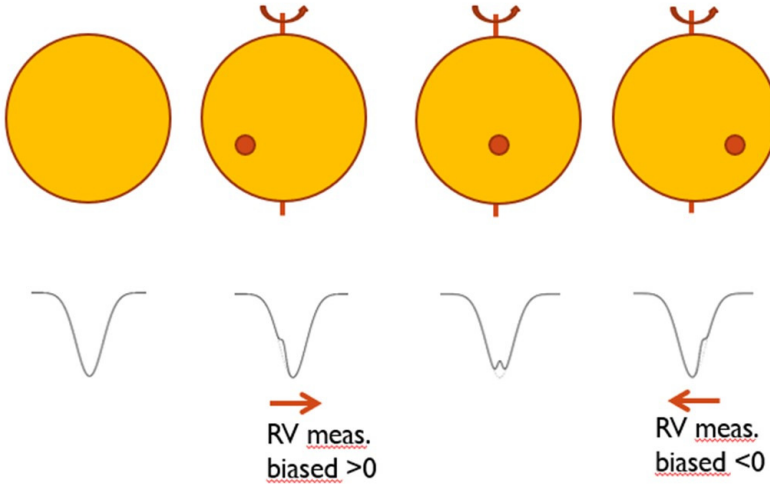


Figure 16: Illustration of the line distortion due to a dark spot and its impact on the RV estimation.

play an important role, although granulation power extends to much lower frequencies, so that the notion of typical time scales is in fact ill-defined. They are followed by supergranulation, with a lifetime of the order of 1-2 days, again with a power spectrum extending to very low frequencies, including in the habitable zone range. Flares are sporadic features, with duration typically below one hour, often down to a few minutes. Spots and plages have a strong effect for periods close to the rotation period, with a complex power spectrum due to the presence of differential rotation and the finite lifetime of these structures, an amplitude which may vary over time (i.e. a modulation of the amplitude on long timescales, for example during the cycle due to the dynamo process, again showing that the notion of timescales is not well defined), and the presence of harmonics. The inhibition of the convective blueshift in plages is also rotationally modulated but has in addition a stronger impact at long timescales (cycle). The gravitational redshift could be due for example to stellar radius variation with the cycle, but the expected amplitude is very small [26]. In the following, I review in more details most of these processes. The Sun will often be used as a reference in this paper. I will provide the corresponding timescales and spatial scales, and describe how they are expected to depend on spectral type.

3.2 Spot and plage contrasts

When rotating across the stellar disk, (dark) spots and (bright) plages change the flux coming from areas of the surface corresponding to different Doppler shifts (due to rotation), i.e. from blueshifted regions to redshifted regions, as illustrated in Fig. 16. As a consequence, the line is distorted at different positions in the spectral line with time (the same principle is used in Doppler imaging or Zeeman-Doppler imaging techniques, when the high rotation rate allows to use this information to reconstruct low resolution maps¹). When measuring the RV from

¹ See for example <http://www.astro.uu.se/~oleg/> and <http://www.astronomy.ohio-state.edu/~johnson.7240/#tomographygallery>.

the line profile (or the CCF), the RV is then biased: for example, in the case of a dark spot, RV varies from a redshift to a blueshift, crossing a velocity of 0 when the spot is on the central meridian. It is reversed for plages (with a slightly different shape due to the position-dependent contrast of plages). When several structures at different longitudes are present and visible at the same time, the contributions from all these spots are therefore partially cancelling each other, producing a complex variability. The resulting RV is therefore a residual between the different contributions, which may lead to degeneracies when attempting to fit this signal.

In the solar case, the root-mean-square (hereafter rms) of the RV is of the order of 0.3-0.4 m/s for both spots [77], plages [105], and the sum of the two (due to partial cancellation). Peaks reach 1-2 m/s when the Sun is very active. The rms of the signal is higher during cycle maximum compared to cycle minimum. The signal is mostly present around the rotation period P_{rot} and at the $P_{\text{rot}}/2$ harmonics. Due to the significant differential rotation and the finite lifetime of structures, there are many peaks around the rotation period, covering a large range in period, and their amplitude strongly varies with time. The bisector also strongly varies with RV. RV are strongly affected by inclination [40], since it is related to the projected rotational velocity, and by wavelength [156], since it is a contribution due to the contrasts. The impact of wavelength is illustrated in Fig. 17. The presence of strong magnetic fields also affects the measurements due to the Zeeman effect [141], which should be more important in IR compared to the optical.

For stars differing from the Sun in spectral types and ages, we expect spot and plage properties to differ from the solar ones. We first consider the properties of individual spots. A number of trends have been identified. First, the ratio between the plage and the spot filling factor (around 10 for the Sun) is expected to depend on age, with spots dominating for young stars [87, 136]. Spot and plage modelling of photometric light curves of young stars also showed a plage-to-spot ratio of the order of unity [e.g. 84]. The spot temperature contrast, which is also an important parameter to determine the impact on the final RV, has been found to increase with increasing stellar effective temperature, from both observations [12] and recent models [132], although they exhibit a large dispersion. The plage contrast is more complex because it depends on μ (cosine of the angle between the normal to the surface and the line-of-sight): the contrast is very low at disk center, and is increasing toward the limb, so that the plage contribution is much larger close to the limb. Models show that the contrast increases toward more massive stars, and also depends on magnetic field [128, 129]. On the other hand, the size and lifetime of stellar spots and plages however is not known. Lifetimes are expected to increase toward less massive stars, because of the lower convection level [leading to a weaker decay see 22, 55, and references therein], and a similar trend has been observed from Kepler data [55]. Furthermore, very long lived spots have been observed on M dwarfs, up to 1-2 yr, for example on GJ674 [17]. This is also the case for younger stars. The size dependence is not constrained however, because of the strong degeneracy between size and contrast, and between spots and plages. There may be some possibilities to derive such trends from in-transit spot modeling of light curves in the future, because the degeneracies are less present [e.g. 36, 151].

Another important consideration is the long term variability of stars, i.e. the modification in terms of activity level over time [for example the presence of cycles or more stochastic variability, see Mount Wilson survey e.g. 9] as well as the average level, related to the number of spots and plages. A useful observable to characterise these properties is the chromospheric emission (see Sect. 2.4). Most publications provide the average activity levels versus spectral type for a large sample of stars [for example 59, 60, 124]. Figure 18 shows recent results [21] for more than 4000 stars, which confirms the results from those earlier publications: for a given spectral type, a very large range of activity levels is covered (e.g. for G stars all

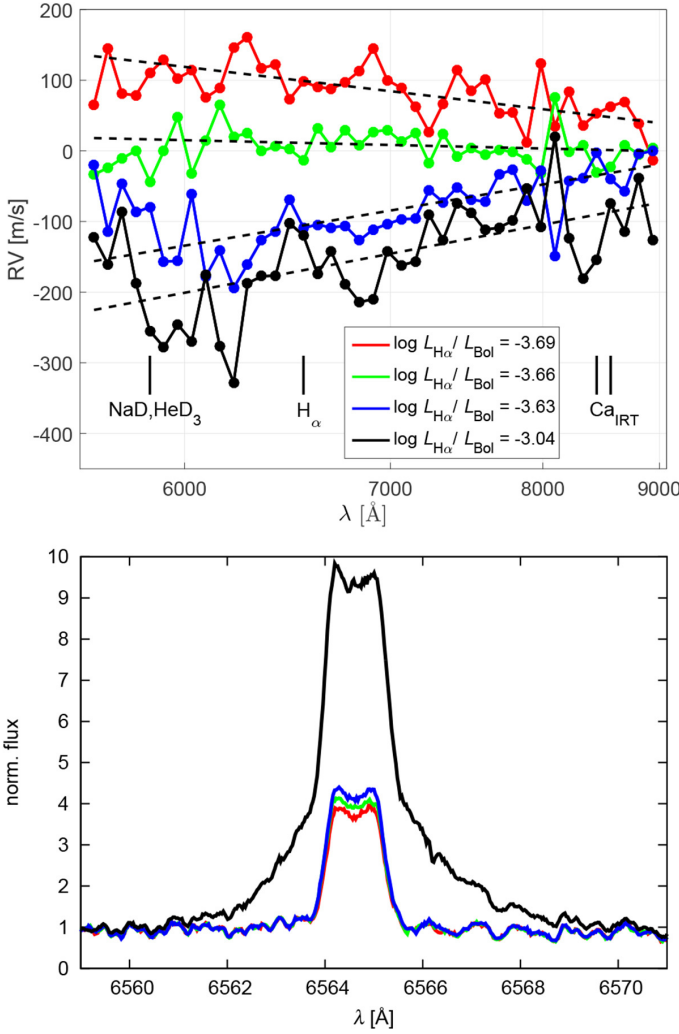


Figure 17: Upper panel: RVs vs. wavelength range (order) for different activity levels (the black one corresponds to a flare), the dashed line indicating the best-fit straight lines to each curve. Lower panel: corresponding $H\alpha$ emission. From [156] reproduced with permission ©ESO.

activity levels are observed), and K stars show a lack of very quiet stars. The cycle amplitude also covers a large range for most spectral types, as shown in [89] for example [see also the adaptation of this relationship in 115], as illustrated in Fig. 19.

3.3 Oscillations

Oscillations such as p-modes affect RV, with many peaks having periods of a few minutes and amplitudes of typically 1 m/s for the Sun. The large amount of peaks defines a well-defined envelope [e.g. 73]. Because of their spectrum, they are easily averaged out [29, 51]. Their amplitude and frequency increase slightly with stellar effective temperature. In addition to

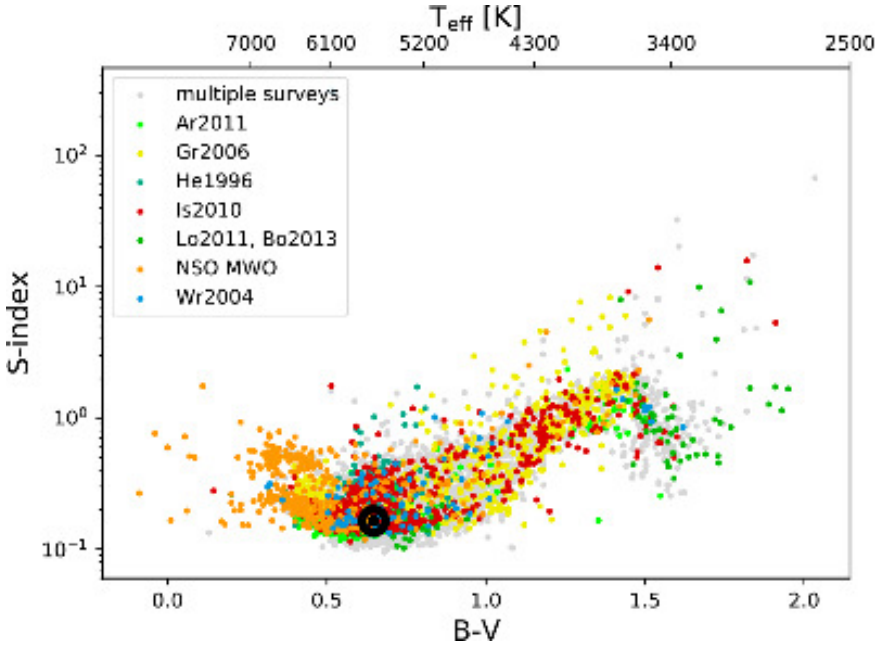


Figure 18: S-index vs. B-V of 4554 main-sequence stars, calibrated to the Mount Wilson scale. The black symbol represents the Sun during cycle minimum. From [21] reproduced with permission ©ESO.

those modes, some specific additional modes could be present: [82] have observed sectoral modes on the Sun, with the main mode having an amplitude of 0.44 m/s and a period of 16.19d (whose value depends on the rotation rate). For other stars, attention should therefore be given to the possibility for such modes to be present when attempting to detect low mass planets in this specific period range. More massive, young stars can show much more intense pulsations, such as δ Scuti and γ Dor, with longer timescales (minutes-hours or more). They can be of very high amplitude (km/s) and therefore critical to detect planets: this was for example the case for the detection of β Pictoris c [80], since β Pictoris exhibits many modes with periods around 30 minutes [74, 101].

3.4 The granulation

Small scale convection (granulation) at the surface of the Sun has been identified and studied for a very long time, while solar images with unprecedented spatial resolution obtained with the recently built Daniel K. Inouye Solar telescope allows to access many details of the cell structures, as illustrated in Fig. 20. Cells have a typical size of the order of 1 Mm and lifetime of 10 minutes, with a very large distribution of the values, leading to about 10^6 cells covering the visible disk. The flows are strong, of the order of 1 km/s (both vertically and horizontally). As a consequence, different realisations of those cells over the disk (i.e. at a given time) differ from one time to the other, so the average velocity is not equal to zero but exhibits a small residual which varies over time and is very stochastic. A recent review on the impact of granulation on RV has been made by [28]. The typical rms (root-mean-square) of these residuals has been found around 0.4 m/s from specific lines [53, 131]. A simulation

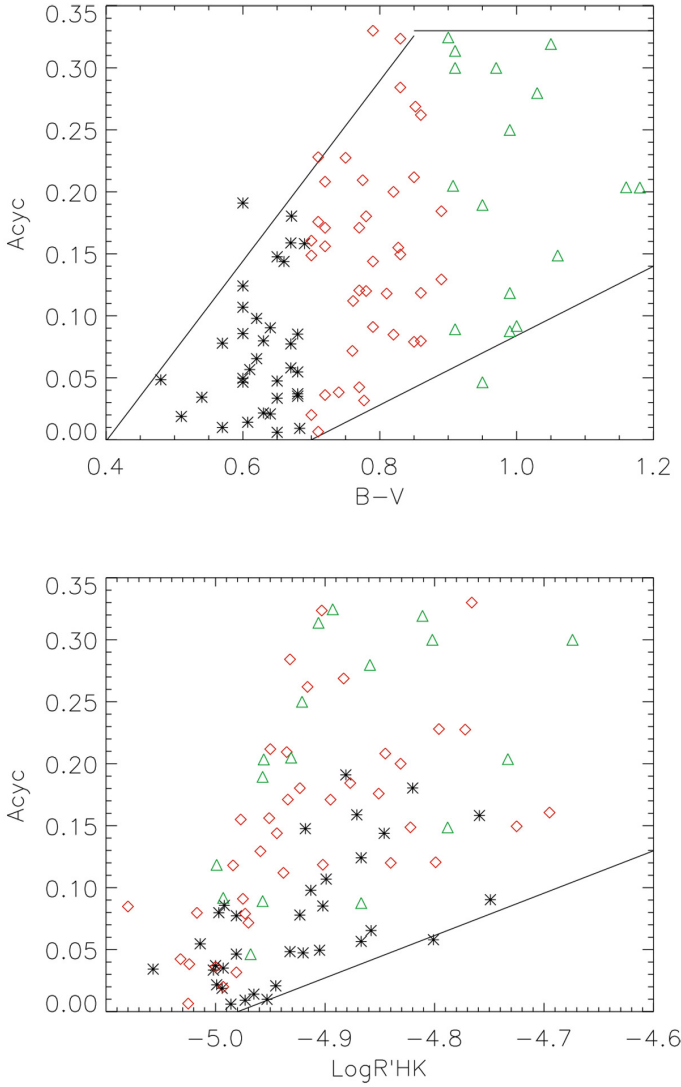


Figure 19: Semi-amplitude of stellar cycles $Acyc$ (in $10^5 R'_{HK}$) vs. $B-V$ (upper panel) and vs. $\log R'_{HK}$ (lower panel), derived from Lovis et al. (2011) after revision of the largest amplitudes, for different types of stars: $B-V < 0.7$ (black stars), $0.7 < B-V < 0.9$ (red squares), and $B-V > 0.9$ (green triangles). From [115] reproduced with permission ©ESO.

made by [114] gave a larger rms (0.8 m/s), while the simulation of [155] based on numerical MHD simulations for the Sodium line as in the observations of [131] provides a rms of 0.4 m/s in agreement with those observations. The simulations of [28] on the other hand leads to a low rms, about 0.1 m/s, which may be due to the choice of magnetic field used in those simulations.

Granulation also strongly affects line shapes from solar observations [e.g. 44] or from simulations [e.g. 11], and the line shape variability can in turn represent a useful indicator of the RV variability [28], although the amplitude seems to be small. The signal can be

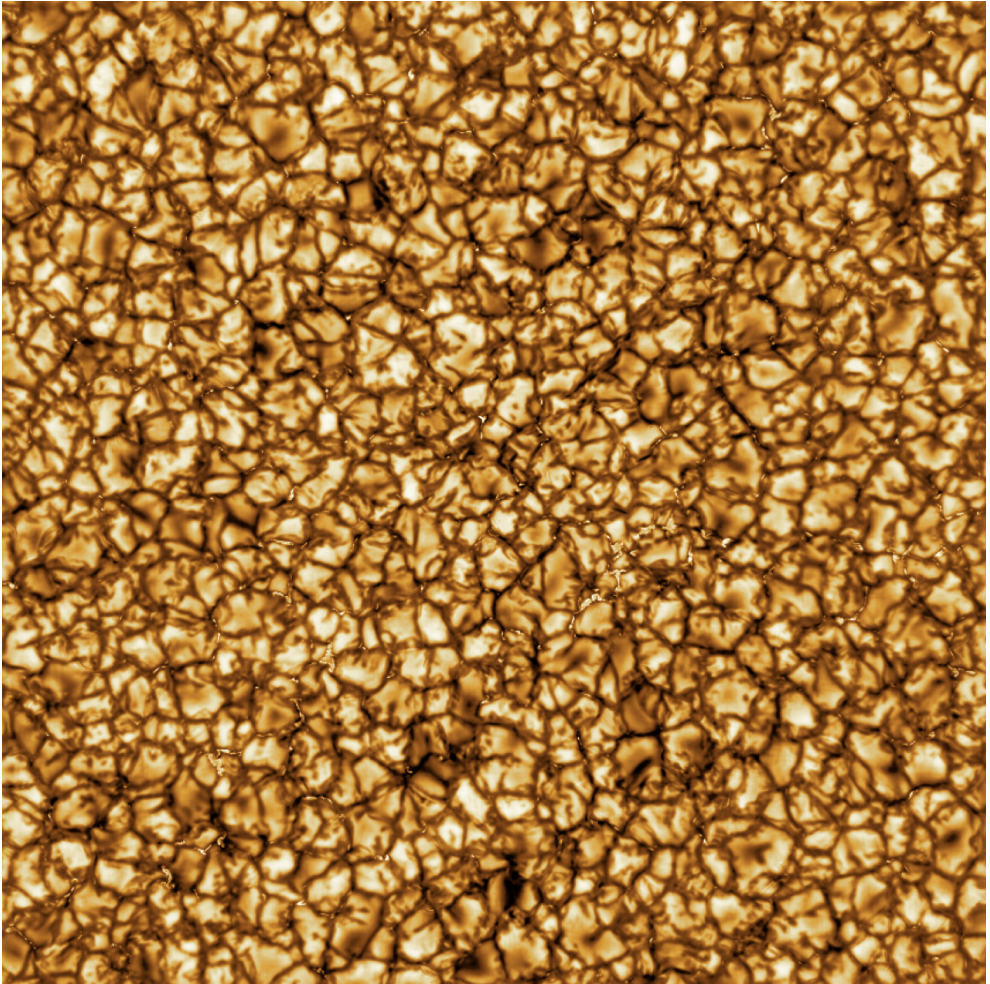


Figure 20: Image of solar granulation from the Daniel K. Inouye Solar Telescope, taken at 789 nm, showing details of about 30 km. Credit NSO/AURA/NSF.

averaged to decrease its rms and its impact on exoplanet detectability [51]. The signal is however very difficult to average out completely [114], because of the specific shape of the power spectrum. The shape of the power spectrum of the granulation contribution has been modeled by [63] by a simple function, with a strong slope at high frequency, and a plateau at low frequency, meaning that they can significantly contribute to the power at all periods of interest for exoplanet searches. This is illustrated in Fig. 21. As a consequence, the rms as a function of the size of the temporal bin reaches an inflexion point around 1 hour, at which point the rms is divided by a factor of about two, then the decrease of the rms RV is extremely slow if we average over longer timescales (Fig. 21). Furthermore, the way the false positive are usually estimated may be biased [153], and [154] proposed to use a standardized periodogram to improve the analysis of this contribution.

Granulation amplitude strongly depends on many stellar parameters, such as temperature and mass, but also log g and metallicity. The velocity field increases with effective temper-

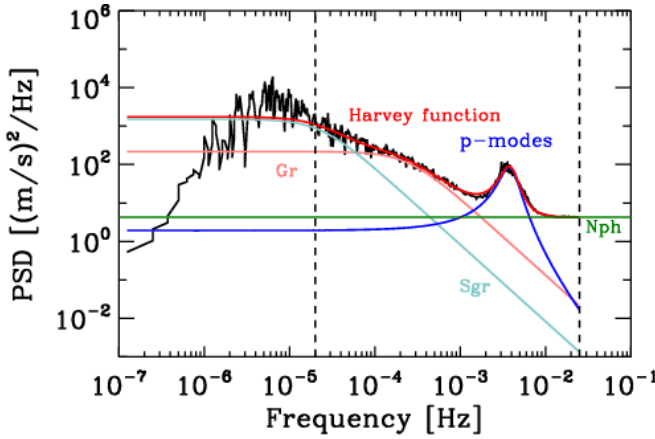


Figure 21: Solar power spectrum from GOLF/SOHO of a subset of ~ 91 days, with a fit of the p-mode envelope (blue), supergranulation (blue-gray), granulation (pink), photon noise (green) and sum of the Harvey function for granulation and supergranulation (red). From [85] reproduced with permission ©ESO.

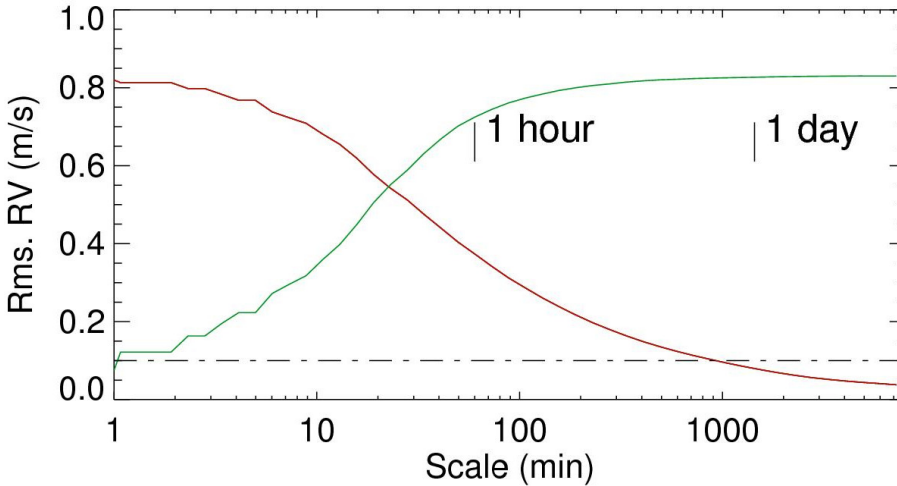


Figure 22: Rms of the smoothed granulation signal (in red) vs. the size of the binning box, while the rms of the residuals are in green. The horizontal dot-dashed line is the level expected from ESPRESSO/VLT. From [114] reproduced with permission ©ESO.

ature from numerical simulations [2, 11, 31, 90, 91, 157, 158]. The same trend is observed from observations, using convective blueshift (see below) estimations [58, 119, 120] and fits of the Harvey's function to observed power spectra for a few stars [51].

3.5 Supergranulation

In addition, flows at a larger scale than granulation are observed on the Sun, called supergranulation, also observed in the photosphere. Reviews on supergranulation can be found in [143]

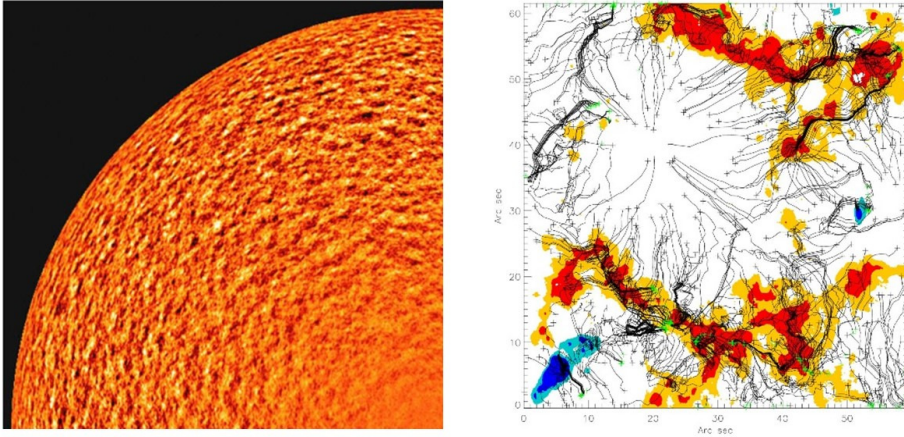


Figure 23: Left panel: Solar dopplergram from MDI/SOHO for a quarter of the solar disk, after averaging and subtraction of the rotation component, to exhibit supergranulation. Right panel: supergranulation flows and network concentration on the edge of a supergranule, from [146] reproduced with permission ©ESO.

and [144]. Supergranules are cells of the order of 30 Mm, with horizontal flows of the order of a few 100 m/s and very low vertical flows (below 30 m/s). These flows can be observed using Dopplergrams, as illustrated on the right panel of Fig. 23. These flows advect the magnetic network toward the edges of the cells (left panel) and can therefore be observed using chromospheric emission maps or magnetograms exhibiting the network, or by tracking granules to reconstruct the flows. The integrated RV is less well constrained than for granulation, but the principle is the same and due to the different realisations of the cell pattern over time. The flows are weaker, but there are also much fewer cells (a few thousands) and therefore we expect the residual to be important as well, while there is no photometric counterpart [e.g. 121]. [114] obtained a range of 0.3-1.2 m/s for the Sun, while [131] observed a rms of 0.78 m/s in a specific line. An example of time series comparing granulation and supergranulation following the Harvey laws are shown in Fig. 21. The resulting signal is much more difficult to average than granulation [114, 115] because of the longer timescale involved: an average over 1 hour does not change the rms at all. [111, 115] showed that supergranulation is more problematic to detect low mass planets (like the Earth) at long orbital periods (habitable zone), because of the amplitude and the low frequency power.

It is difficult to quantify how supergranulation varies with spectral type, especially since the origin of supergranulation is not well understood. [51] performed fits of the Harvey function for a few stars observed by HARPS, but the sample is too small to derive a trend. [113] scaled it to granulation amplitude because [146] showed that supergranulation appears to be related to the behavior of granulation (through the analysis of trees of fragmenting granules).

3.6 Convective blueshift inhibition in plage

Another effect of the presence of granulation is the convective blueshift [e.g. 44], due to the different contributions to the integrated spectrum over the surface coming from the bright up-flow areas of granules (large fraction of the surface, blueshifted) and from the dark downflow

areas (small fraction of the surface, redshifted). The amplitude of this convective blueshift for the Sun is about 400 m/s [see 140, for a recent estimation]. If the convective pattern is attenuated (with smaller weaker granules) by a different amount over time, for example in magnetic plages, the resulting convective blueshift is modulated, with a net redshift at time when there are more plages compared to a quiet star. This leads to a modulation of the RV at the rotational period (as plages cross the disk) and on the long-term (during the activity cycle for example, since the signal depends on the amount of surface covered by plages). The simulations of [105] showed that the resulting signal for the Sun over a complete cycle is the dominant contribution to the RV variation, with a long term amplitude of about 8 m/s. It also contributes significantly to the rotational modulation. The periodogram of the time series shows that it is responsible for a strong peak at the cycle period as expected, but also of large power in the habitable zone, which is significantly higher than the power due to a low mass planet like the Earth. For the Sun, [105] showed that the amplitude of the convective blueshift attenuation was much stronger in large plages (where the magnetic flux is higher) compared to smaller plages and small network features, with a ratio of about 6 between the maximum and the minimum [and furthermore small features show a weaker cycle dependence, 103]. Such a trend was also observed later by [123].

The convective blueshift has interesting properties which can be used to quantify this effect in other stars and also to develop mitigating technique. The most important one is that it depends on the line depth, with deep lines (formed higher in the atmosphere) exhibiting a much smaller convective blueshift than weak lines, formed higher in the atmosphere [e.g. 44, 140]. This property has been exploited by [112] and [46] to propose different mitigating techniques based on spectral line selection (See Sect. 5.1).

Several applications of this property have been made in various papers to stars other than the Sun [3, 42, 43, 81]. [58] used this property to compare the convective properties of different stars: he computed the position (which is then converted into a velocity) of the bottom of selected lines in a 100 Å range as a function of the line depth and deduced that all stars showed a similar shape, i.e. a universal signature, which he called the third signature. Following this work, [119] and [120] assumed a direct relation between the shape of the signal and the amplitude of the convective blueshift for a large sample of spectral lines and a large sample of stars: this allowed to confirm the trend of the amplitude of the convective blueshift with spectral type. It also allowed to observe a dependence of the convective blueshift on the average activity level (for a given spectral type), showing that it was indeed weaker in active stars: an attenuation factor (characterising the attenuation of the convective blueshift in plages) could be deduced from this analysis, with a possible trend with spectral type between 5400 and 6400 K.

3.7 Evershed flows

On the Sun, outward horizontal flows have been observed around sunspots and characterised for a long time (upper left image in Fig. 15). These flows are of the order of 1-2 km/s. If they are symmetric, a very small residual is expected after integration on the spot, but the resulting RV could be slightly higher, especially for large irregular spots. The effect has however not been quantified precisely [see 66, for discussions about photospheric flows associated to spots and active regions].

3.8 Meridional circulation

Meridional circulation constitutes a contribution to the integrated RV which has not attracted much attention compared to other processes. It is a large-scale flow which, if variable with

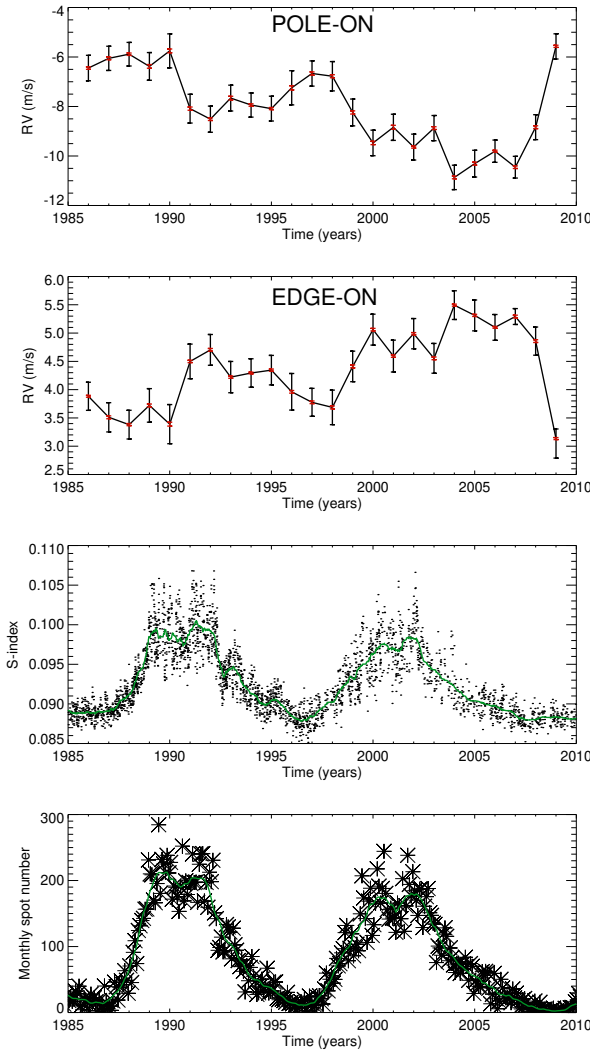


Figure 24: Solar integrated radial velocity due to meridional circulation reconstructed from the observed latitudinal profile of [161], over two solar cycles, for pole-on configuration (first panel), edge-on configuration (second panel), compared to the activity cycle from the chromospheric emission S-index (third panel) and spot number (fourth panel). From [110] reproduced with permission ©ESO.

time, should also lead to RV variability after integration on the disk. On the Sun, meridional circulation is a poleward flow, with a maximum amplitude in the 10-20 m/s range and varying with the solar cycle. It has been studied for several decades using various complementary techniques, such as Dopplergrams, magnetic feature tracking, or local helioseismology [52, 64, 70, 75, 76, 102, 127, 152, 162]. Converging flow patterns have also been observed in activity belts [86, 102, 164]. The first study estimating its impact on integrated RV has been made by [93], who performed a reconstruction of the solar variation, for the Sun seen

edge-on: the resulting contribution was however superimposed to other large contributions and therefore its precise behavior was not clear. [110] have therefore performed a new reconstruction, based on the solar measurements made by [161] covering two solar cycles and for different inclinations between edge-on and pole-on. The resulting amplitude shows a significant variation with an amplitude over cycle 23 of the order of 2 m/s for the edge-on configuration and a few m/s for the Sun seen pole-on, with a reversal in sign between the two orientations for an inclination around 50° . The results are shown in Fig. 24. The impact on exoplanet detectability is therefore important, especially for the extreme configurations, while intermediate inclinations may be the most suitable in that respect.

For other stars, the amplitude of meridional flows is decreasing for fast rotators [10, 23] and for low mass stars [23, 96]. [110] extrapolated the solar amplitudes of cycle 23 to a range of F6-K4 old main-sequence stars of various activity levels using the scaling laws from [23] and assuming a proportionality to the cycle amplitude. They found an amplitude between 0.1 m/s for the most quiet stars and intermediate inclinations, up to 4 m/s for the most active stars. The amplitude should be much smaller than those results for fast rotators if they exhibit a multicell pattern as expected from HD simulations [61, 62, 96] however.

3.9 Flares

Flares produce a RV signal which is very stochastic and of very short duration. Because they are very localized on the surface and represent a small area, their impact is negligible for Sun-like stars. M dwarfs RV time series are often affected by flares however, because some of these stars are extremely active and exhibit very energetic flares: they appear as “outliers” superimposed to the RV time series with amplitude of up to a few 100 m/s.

4 Approaches to the problem

As shown in the previous section, there are many sources of stellar variability impacting RV at various timescales. Many of them are in the range 0.3-1 m/s for the Sun, with a complex behavior for solar type stars, while the activity patterns may be more stable for young stars or M dwarfs. Such a complexity is due to the activity pattern itself, the evolution and finite lifetime of the structures, differential rotation, degeneracies between some contributions and complex temporal variability. None of those contributions are strictly periodic. In practice, the analysis of RV time series is also affected by the temporal sampling since it is most of the time sparse and irregular. To deal with this complexity, several complementary approaches have been developed by many groups. I will first briefly review the mitigating techniques which have been implemented. Then I will discuss the approaches used to understand and quantify the contribution of the different processes to the RV measurements, and to test and improve mitigating techniques.

4.1 Mitigating techniques

Given the very serious challenge represented by the presence of the stellar activity contribution to the RV time series when searching for exoplanets, many techniques have been implemented over the last 20 years. They all contribute to remove part of the stellar signal although none of them so far allows to reach very low levels of residuals, although a systematic analysis of the impact and performance is seldomly made. We list these approaches in Table 1.

Table 1: Summary of mitigating techniques. References indicate papers proposing the method or using it, but the list is not always exhaustive. For example, many analysis rely on the use of Gaussian processes and chromospheric emission indicators. (*) including convective blueshift inhibition in plages at the rotational timescale (**) including convective blueshift inhibition in plages at all timescales (***) including convective blueshift inhibition in plages at long timescales.

Methods based on usual RV time series only		
Fits of sinusoids around P_{rot} and harmonics	spot and plage (*)	[14]
Prewhitening of the signal at P_{rot}	spot and plage (*)	[134]
Spot modeling	spots and plages (*)	[65] [125] [47] [69]
Averaging of the signal	granulation	[51]
Periodogram standardization	granulation	[154]
Methods based on activity indicators computed from the spectra		
Correlation with the bisector span	spot and plages	[135] [40] [15]
Correlation with the chromospheric emission	plages (**)	[15] [50] [106] [145] [137] [83] [19]
Use of gaussian processes based on those activity indicators	spots and plages (*)	[137] [48] [94] [37]
PCA analysis	spots and plages	[38]
Doppler imaging techniques	spots	[68]
Methods based on RV time series computed from subsets of data		
Produce independent time series based on different sets of lines	plages (***)	[112]
Use of selected lines with different sensitivity to magnetic field	spots and plages (**)	[46] [35]
Wavelength dependence of the signal and chromatic index	spots and plages	[156]
Methods based on activity indicators obtained separately from the spectra		
Estimation of the RV signal using the photometric light curve (FF' method)	spots	[1]

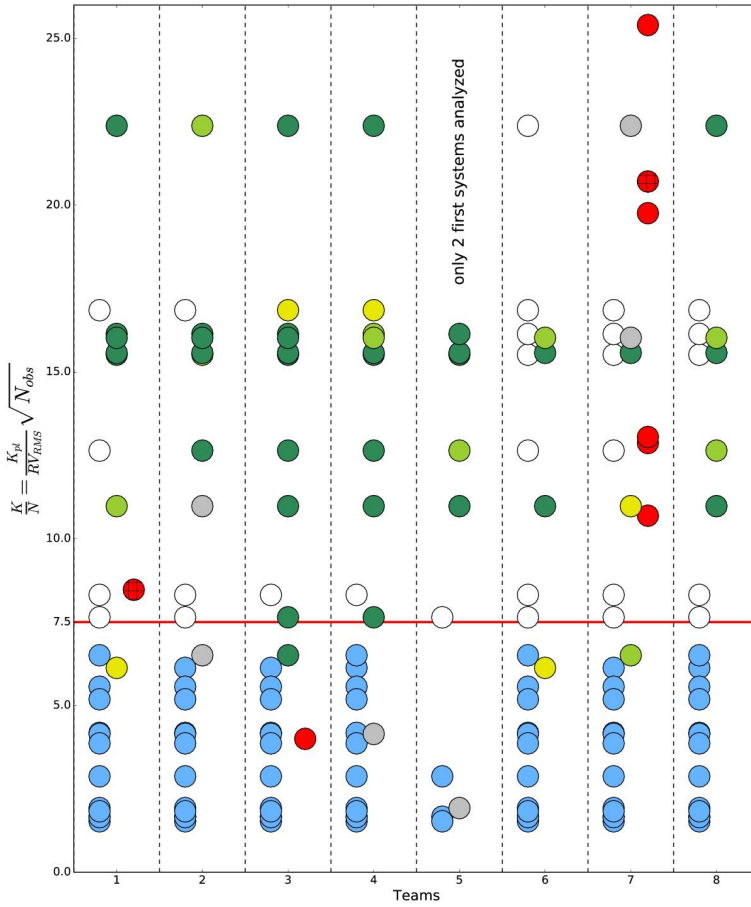


Figure 25: Results from the fitting challenge organized by [48], representing K/N (see text) for each team and all planets in five synthetic systems. The color flags indicate the status of the detection (or lack of detection): planet not detected ($K/N < 7.5$ in blue, $K/N > 7.5$ in white), planet properly detected (green), detected planet but wrong parameters (yellow if claimed, gray if probable), false positive or negative (red), probable detection but no planet injected (orange). From [48] reproduced with permission ©ESO.

All these methods significantly reduce the stellar contribution to some level, but they never remove it completely. However, they also all have limitations and lead to crucial questions. First, since none of them removes completely the stellar signal, what is their limit in terms of performance? Second, how reliable are the residuals: do they remove part of any planetary signal, do they add any spurious signal? There are many reasons for these limitations:

- Strong degeneracies between the numerous processes are present, which make them difficult to estimate.
- Involved processes are very stochastic, with a very complex frequency behavior of the stellar processes, and contributions at all timescales.

- The models used in these methods are not perfect and do not describe with enough accuracy the processes (or do not include any physics at all), and do not take all processes into account.
- Some parameters needed in those models are not always properly estimated, for example the rotation period, which cannot be uniquely defined for a given star because of differential rotation.

Finally, the very sparse sampling and the presence of photon noise complicate the situation. In addition, very recent results by [147] also show the possible contamination of RV measurement above the 10 cm/s level due to the Moon (solar light). We also note that in principle, some of these methods should provide interesting insights on stellar activity itself (such as the rotation period, properties about spots and plages, and on activity cycles), although they usually focus on exoplanets only and therefore on the residuals after subtraction of part of the activity contribution.

The fitting challenge organized by [45] allowed to test some of these methods (8 teams participated to the challenge) on a few complex time series, most of them synthetic (see next section) in a blind test. The results are presented and discussed in [48]. The injected planets covered a very large range in period (~ 0.8 -3400 days). Some of these methods allowed the teams to retrieve some planets well (mostly the techniques based on Gaussian processes) while other did poorly (no planet or wrong planet retrieval). The same was true for the retrieval of the rotation period. They also propose a criterion defined as $K/N = K_{\text{pla}} \sqrt{N_{\text{obs}}}/\sigma$, where K_{pla} is the amplitude of the RV planet signal (which can be converted into a mass, providing an orbital period) and σ is the RV jitter after a correction of the RV time series using a linear correction with $\log R'_{HK}$ and second degree in time polynomial fit. As shown in Fig. 25, they evaluate an approximate value of 7.5 for this criterion, above which the retrieval performance was bad and below which the performance was good in many cases. This shows that given current techniques, the retrieval of low mass planets (such as the Earth) in the habitable zone around solar type stars is not possible [as also shown by 107, on a large set of activity simulations using that criterion, see next section].

4.2 Simulations

An important way to study in more detail the stellar processes affecting RV measurements and their impact are simulations. I define here three categories. Simulations involving a very simple magnetic configuration (typically 1-2 spots) for a given set of parameters and simulations involving complex activity (spot, plage) patterns aiming at reproducing solar-like activity time series, have been implemented to reproduce the impact of magnetic activity. The general approach for these two categories is the following, although some of the steps can be omitted in simplified versions: at each time step, structures (spots, plages) are defined (position, size, contrast), localized on a map or not, with spectra associated to each contribution from the surface, providing a final spectrum after integrating on the disk. This spectrum can then be analysed like an observed spectrum (after adding white noise for example, or extracting a certain sampling). A third category concerns photospheric flows. I review their objective and a few selected results below.

4.2.1 Simulations with 1-2 spots

The objective of simple simulations, i.e. with typically 1-2 spots, is to derive typical amplitudes and shapes for simple activity configurations, in order to understand and identify fine effects related to a single structure at the rotational timescale. When the star exhibits a simple

configuration, it can also be used to fit actual observations. Different tools have been implemented independently (list not exhaustive), such as SAFIR [40], SOAP and SOAP2 [13, 47], and Starsim [69] for this purpose.

I list here a few representative results. [40] simulated a single dark spot (plage and convective blueshift inhibition were included later) and studied the impact of $v \sin i$, inclination, spectral type, and spot position, as well as the relationship between RV and BIS (see Sect. 2.4). The main results are the following: the BIS does not exhibit any temporal variation if the spectral resolution is low; there are regimes with significant stellar RV variations but no BIS variations; the general behavior can be described as scaling laws depending on the instrument and spectral types; latitude and inclination strongly impacts the signal; and chromatic effects are observed as expected. They note that data inversion should be difficult to perform because there are too many parameters. [13] obtained similar results, but also studied the effect of latitude in more detail as well as the impact of the limb darkening function. They also performed a comparison with observations. Their model was improved in [47], with the addition of plages and in particular the inhibition of the convective blueshift. The spectra of the quiet Sun and of spots were used as inputs and they provide detailed temporal variations depending on various assumptions for those spectra. As an illustration of an application of the tool, the modeling is presented for two stars, one where a single plage is dominating (α Cen B, plage size of 2.4% at a latitude of 44° and an inclination of 22°) and one where a single spot is dominating (HD189733, spot size of 0.8% at a latitude of 61° and an inclination of 80°).

4.2.2 Simulations with complex activity patterns

The objective of complex activity patterns is also to identify and study fine effects, but for more complex configurations at the rotation period as well as on long timescales (for example taking the variability of the activity level into account). It also allows to study detectability and characterization performance for solar-like stars using a systematic analysis (such as the impact of different samplings, tests of correction methods...) and to find new diagnosis to develop new correction techniques. This has been made for the observed Sun [67, 77, 93, 105, 106, 117], for a simulated Sun [20] and then for other stars: for a few stars in the fitting challenge [45], for spots only [149] and for a large range of stellar parameters in a systematic way [107, 115, 116]. Results from the fitting challenge are presented above (Sect. 4.1). Here I focus on the approach described in this latter series of papers.

The best characterised star is the Sun, and the manifestation of solar activity has been the subject of a huge effort from the solar community for decades. I first focus here on the manifestation of magnetic activity and in particular the contribution directly due to spots and plages (described in Sect. 3.2 and 3.6). The starting point of the approach was the question: if we were to observe the Sun from the outside, would we be able to detect the Earth given our current instruments and methods? The study focused on the system star-planet (or Earth-Sun) without the presence of additional more massive planets like Jupiter which would have to be taken into account for a complete solar system and which would be responsible for additional “noise”. The main principles of the approach were: 1/ use of our knowledge of the Sun and extending the analysis to other stars; 2/ simulations of complex activity patterns; 3/ systematic approach in terms of parameter space and performance evaluation; 4/ use of these simulations to find new diagnosis and improve correction methods.

A first step was performed by reconstructing the solar integrated RV due to magnetic activity using actual observed structures (spots, plages) and a model [77, 105]: this allowed to study the Sun seen edge-on during cycle 23 and showed that the dominant contribution came from the inhibition of the convective blueshift in plages, with a long-term amplitude

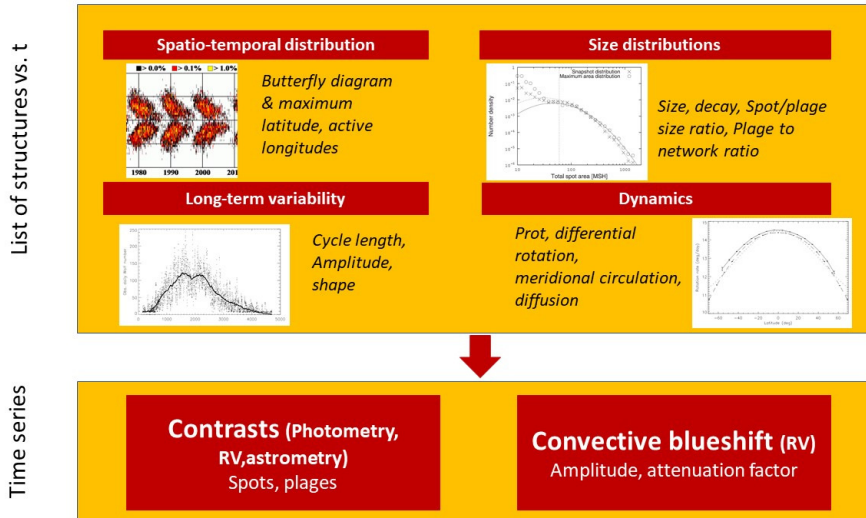


Figure 26: Principle of the simulations performed in [20] for the Sun. The first step consists in generating spots and plagues (position, size) at each time step based on various categories of empirical laws listed here. The second step generates the time series of observables.

of the order of 8 m/s, i.e. two orders of magnitude higher than the Earth signal. This was illustrated in Fig. 14. This was confirmed with a reconstruction of the solar integrated RV from MDI/SOHO Dopplergrams [118], as well as from HMI/SDO Dopplergrams [67]: this is presented in Sect. 4.4. A model (Fig. 26) was then implemented to produce realistic sets of spots and plagues in the solar case, based on empirical laws such as the solar butterfly diagram, the cycle shape, size distribution, plague-to-spot ratio distribution etc. [20]. RV time series were then computed from the list of structures as in the reconstruction made in 2010. This model allowed to study the dependence of the variability on inclination. At low inclinations compared to the edge-on configuration, the amplitude of the signal around the rotation period strongly decreases, although there is still some contribution present around similar scales due to the evolution and lifetime of structures. The long-term amplitude slightly decreases (due to the fact that plagues, at low latitudes, cover a lower apparent proportion of the disk), but is still important (a few m/s). This step also allowed to be able to extrapolate this work to solar type stars, which was done in [115], described below.

In [115], the extrapolation to solar-like stars was limited to F6-K4 stars and old main sequence stars [plage dominated regime 87] to select a regime where some of the solar parameters would still be reasonable (Fig. 27). In that domain, for each activity level and spectral type, a number of parameters are explored to cover various configurations. Rotation rates for example depend on activity level and spectral type [95] while the amplitude of the convective blueshift inhibition can be deduced from the analysis of a large sample of stars [119, 120]. A good knowledge of stellar activity (observational and theoretical) is necessary to perform such simulations. The parameters which are explored within a range compatible with observations for each point in Fig. 27 are: rotation period, cycle period and length, spot contrast, and maximum latitude of the butterfly diagram. All laws are described and discussed in [115]. Other observables than RV were also produced: chromospheric emission $\log R'_{HK}$,

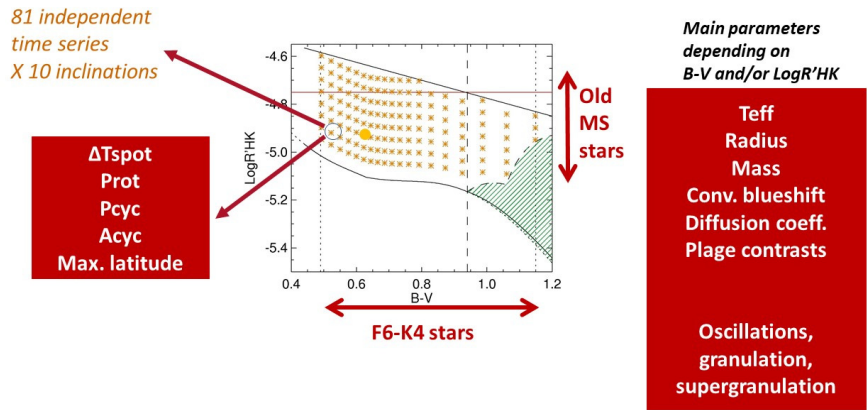


Figure 27: Range of stars covered by the simulations of [115], extrapolating the solar simulation of [20] described in Fig. 26 to other stars. Parameters adapted to different stellar types and activity levels are indicated.

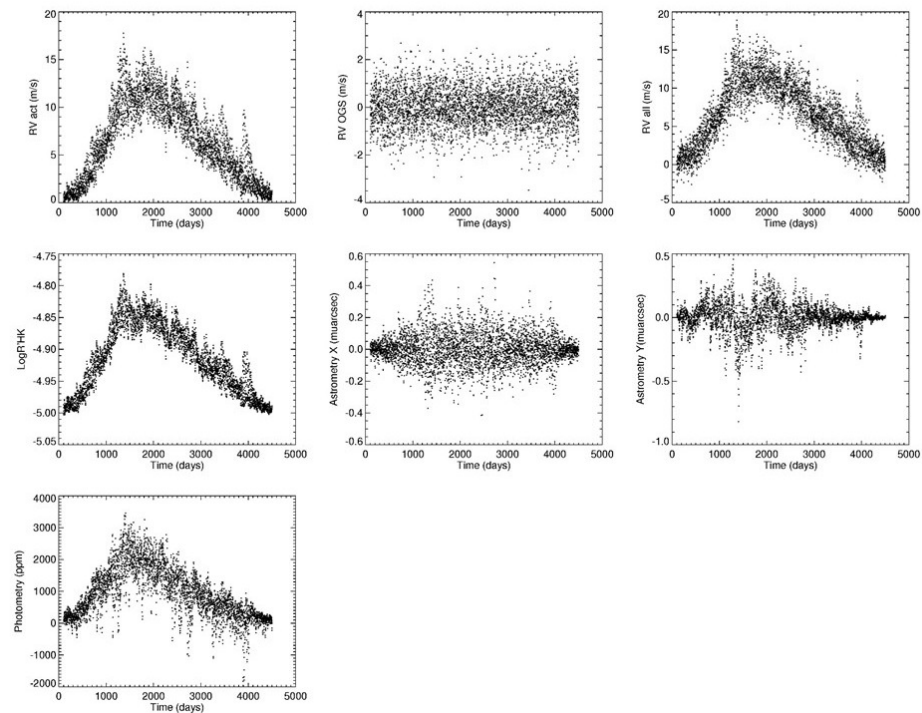


Figure 28: Example of time series produced by the simulations made in [115].

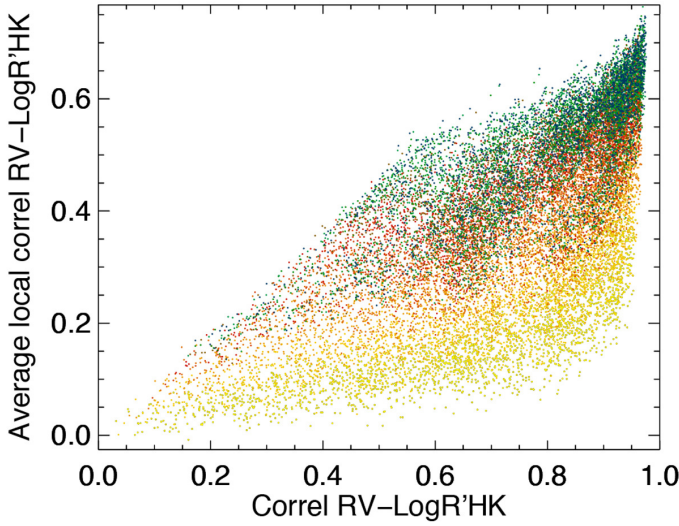


Figure 29: Average local correlation between RV and $\log R'_{HK}$ vs. the global correlation over a large sample of simulations. The color code indicates stellar inclination (from pole-on in yellow to edge-on in blue). From [116] reproduced with permission ©ESO.

photometry [108], and astrometry [113]. An example of time series is shown in Fig. 28 (one realisation, edge-on configuration).

The properties of the synthetic time series have been compared to other variables, for example photometry, and $\log R'_{HK}$, and to published RV jitter trends with spectral type [163, for example] in [107]. They show a good agreement of the trend, although the RV jitter is smaller in the simulations. In particular, they discussed the fact that given our current understanding of stellar activity, it is impossible to reproduce a jitter of a few m/s in the case of very quiet ($\log R'_{HK}$ lower than -5 for example) stars, meaning that either the uncertainties has been underestimated in published RV jitters, or that an additional contribution (not stellar, for example planets, or stellar, such as meridional circulation) are present. However, there is a good agreement of the RV versus $\log R'_{HK}$ slope with the observations of [89]. In addition, Fig. 29 shows a comparison between the average local correlation (computation on small subsets of the synthetic time series, typically at the rotational timescale) between RV and $\log R'_{HK}$, and the global one (computed over complete time series, i.e. dominated by the cycle of activity). This shows that although the long-term correlation is often close to 1, the local correlation can be much smaller. It also strongly varies for a given time series depending on the current activity level for a given time series.

A simple analysis of the RV jitters allows to estimate the order of magnitude of the detection limits for this range of stars. Using the criterion proposed by [48] (see above for definition, $K/N \sim 7.5$), a rough estimate of the typical mass which could be detected with current techniques can be performed. [107] show that with a low number of observations (100 nights, representative of current ordinary time series), the minimum mass is of several M_{Earth} , and only for low mass stars (the threshold is above 10 M_{Earth} for F stars), with maximum values of several tens of M_{Earth} , for planets in the habitable zone. A 1 M_{Earth} level can only be reached for a very large number of nights (several thousands over a long period) and low mass stars (K stars mostly). A more sophisticated approach will be the subject of a future work.

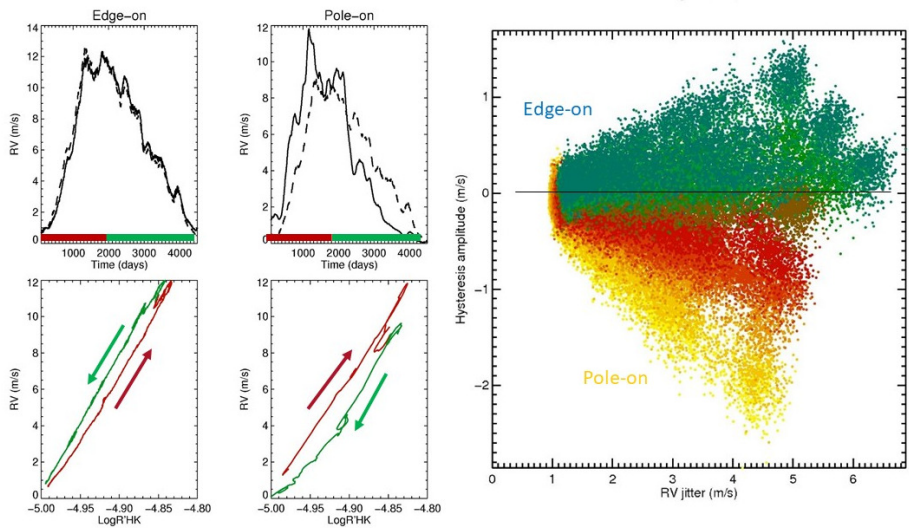


Figure 30: Left panels: example of smoothed synthetic time series (RV as a solid line and scaled $\log R'_{HK}$ as a dashed line, upper panels) and RV vs. $\log R'_{HK}$ (lower panels). The ascending phase of the cycle is in red and the descending phase in green and have a different behavior, as shown in [116]. Right panel: amplitude of the hysteresis vs. RV jitter for different inclinations (from pole-on in yellow to edge-on in blue), from [116] reproduced with permission ©ESO.

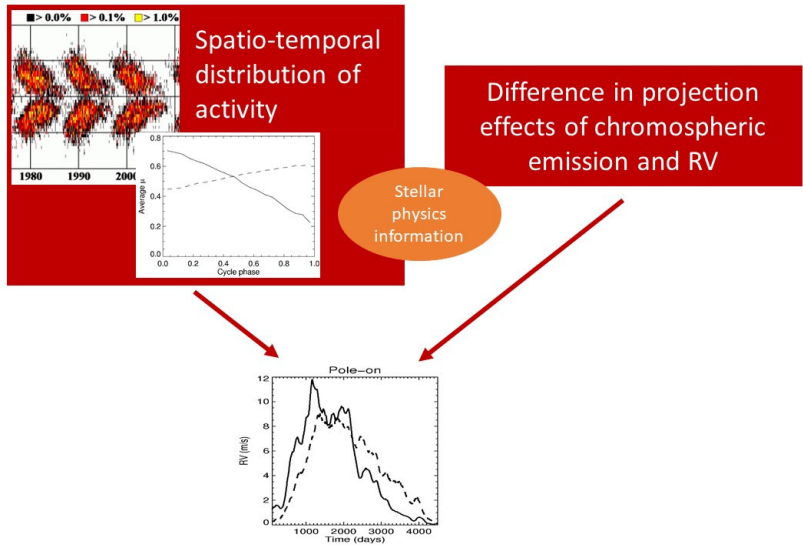


Figure 31: Physical processes leading to the hysteresis pattern observed in Fig. 30. The butterfly diagram (NASA/NSSTC/HATHAWAY) leads to a change in average μ with time, which depends on inclination. The average μ vs. cycle phase and RV vs. time are from [116] reproduced with permission ©ESO..

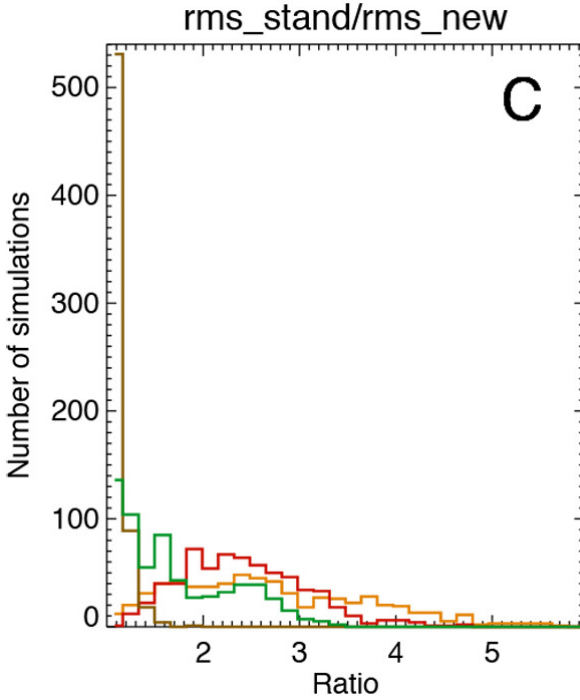


Figure 32: Gain in rms (on the smooth synthetic time series) for four different inclinations, G2 stars, with a standard correction method (linear correlation in $\log R'_{HK}$) and the new method taking the combined effects of the butterfly diagram and projection effects into account. From [116] reproduced with permission ©ESO.

Finally, [116] have shown that this type of approach can be used to improve correction techniques. A widely used method to remove part of the stellar signal is to apply a linear correlation between RV and $\log R'_{HK}$ [e.g. 15, 19, 41, 50, 83, 137, 145]. It is well known that it is not perfect [e.g. 117], but it allows to remove a significant part of the signal when the inhibition of the convective blueshift is dominant, because both are strongly correlated with the plage filling factor. These simulations put a departure from the linear correlation into evidence, and are useful to understand why there is such a departure and improve the correction techniques. This is illustrated in Fig. 30: the long-term variation shows a different relationship during the ascending phase of the cycle and during the descending phase. This is due to two effects related to the geometry, the dynamo processes and surface processes (Fig. 31): 1/ the butterfly diagram (activity pattern as a function of time and latitude, shown in the upper left panel of Fig. 31) leads to a different average μ over time: a decrease when the star is seen pole-on and an increase (to a lesser extent) when the star is seen edge-on, because the activity pattern moves from mid-latitudes to the equator during the cycle; 2/ RV and chromospheric emission suffer from different projection effects. This leads to a distortion of the RV time series with respect to the $\log R'_{HK}$ time series, which could be corrected by using a correlation with the $\log R'_{HK}$ times a function describing the contribution of these two effects. The application of this technique led to a significant improvement of the long-term residuals in those simulations (Fig. 32).

4.2.3 Simulations of the small or large scale dynamics

The simulation of photospheric flows has taken different forms in the literature. For meridional circulation, it has been performed using direct observations of the flows, i.e. the observed latitudinal profiles over time [93, 110]. For granulation and supergranulation, some simulations have also been based on properties deduced from hydrodynamical simulations for granulation and/or supergranulation by [114] using the simulations of [142], directly from those hydrodynamical simulations [24, 25, 27, 155], and from the Harvey's law [109, 111].

For stellar meridional circulation [110] extrapolated the solar results using the results of [23] from MHD simulations. Some results have been presented in Sect. 3.4, 3.5 and 3.8 and are not repeated here. This type of approach allows to model synthetic RV time series, which can be used to characterize the amplitude of the contributions due to these processes, but also, as for magnetic activity, to test performance (detection rates, false positives) and to improve correction techniques.

4.3 Stellar observations: simultaneous campaigns

Another way to understand better RV time series is to perform simultaneous RV observations with other observables. Spectroscopic indicators as described above (such as line shape indicators and $\log R'_{HK}$ measurements) are widely used, but it is important to note that they are either noisy (bisector shape for example) or representative of one of the contributions only: chromospheric emission is mostly sensitive to plagues and in the end representative of the convective blueshift inhibition contribution. The use of simultaneous photometry, which is sensitive to the contribution to the contrast of spots and plagues is therefore particularly interesting to constrain better the different processes at play, although it does not model all of them. [1] has for example proposed the FF⁺ method, to reproduce the RV signal due to spots (mostly for 1-2 spot configurations) from the photometry. Simultaneous measurements have therefore been done [e.g. 33, 88] for a few stars. Such coordinated campaigns, aiming at covering well the rotation period for a good sampling, are very difficult to implement, because ground-based photometric observations are much noisier than from space and are subject to meteorological conditions, but they will be useful in the future to provide a complementary approach. The model derived from photometry is also incomplete due to the presence of other process in RV, as described in Sect. 3 [e.g. 67].

4.4 Observations of the Sun as a star

Finally, as already pointed out in previous sections, our knowledge of the Sun is crucial to progress on these issues. Traditionally, there have been a lack of long-term stable integrated RV for the Sun, since most studies focused on observations with a good spatial resolution. [100] obtained an upper limit of 4 m/s using deep lines (which are not very sensitive to the convective blueshift contribution at small wavelengths), while [39] observed a peak-to-peak amplitude of 30 m/s in the infrared (2.3 μm) and [72] obtained a large variability (30 m/s on the long-term, 20 m/s on the short-term) using the K 7699 Å line. Later on, the reconstruction of the solar RV using Dopplergrams from MDI/SOHO by [118], whose objective was to reproduce the contribution of active regions (and mostly the convective blueshift inhibition due to the noise level), was compatible with simulations as described above, with a long-term amplitude of the order of 8 m/s, which confirmed the reconstructions made from observed structures by [105] and later by [67] from HMI/SDO. Other solar observations were then performed using an indirect approach by [83] (observations of Jupiter satellites, asteroids, and the Moon), with a variability compatible with those previous results.

A significant step has been made with the advent of “Sun as a star” observations using stellar high-performance stabilized spectrograph, the solar light being supplied to the spectrograph through a coelostat and an integrated sphere. Even though this reproduces stellar observations, some adaptation in the data processing had also to be done due to the finite solar apparent diameter, which led to some spurious effects (differential effects due to the atmosphere for example). Such a device has been implemented on HARPS-N for the last 4 years [34, 49, 133]: the Sun is observed about 6 hours per day with a 5 minute cadence. It has also been implemented more recently on HARPS in La Silla, or on Expres (Lowell observatory) and several similar projects are on-going. Some results have already been obtained by [123]. The comparison with solar reconstruction from Dopplergrams and other indicators such as the unsigned magnetic field is also very promising [66].

5 Conclusion

Stellar activity contributes to RV in a complex manner because of the large number of processes, covering a large range of time scales but also similar orders of magnitude. There is also a large amount of degeneracies between the different contributions. A large diversity is observed among stars of different spectral types and ages. It is however complementary to other observables (in particular because some contributions can be seen only in RV) and it is also crucial to understand stellar activity observed in RV given the challenge it constitutes to observe low mass planets. Techniques to push the limits to be able to detect low mass planets (like the Earth), and to characterise their mass, including at long orbital periods (habitable zone), will therefore have to be improved and be based on our physical knowledge of stellar activity to be able to control the residuals after correction. Recent studies also showed that there was still to learn about the Sun to fully understand its integrated RV. The development of highly performant instruments such as Espresso is critical to improve our understanding, because they allow to develop more sophisticated techniques, for example dependent on spectral lines for example (requesting high S/N even after computation on a subset of lines). A large wavelength coverage in the optical and infrared is also extremely promising.

One important conclusion however if we want to be able to push the limits toward very low mass planets in the habitable zone is the fact that a huge number of points (> 1000) will be necessary to better take stellar activity into account, which could constrain future facilities. The fitting challenge organized by X. Dumusque [45, 48] or the complex activity pattern simulations [115, and following papers] also show that given the challenge and the complexity of stellar activity (especially its high level of stochasticity and the lack of independent activity indicators for certain contributions), it will be necessary to develop new techniques, and probably to combine them to be able to obtain robust results in terms of detection.

In this context, it is also important to recall the other factors contributing to the difficulties of studying stellar activity in RV, in particular the usually very sparse and irregular sampling (apart for the Sun) and the fact that it could be superimposed to other contributions, and in particular undetected planets (or massive planets removed but with some uncertainties on the parameters): stellar activity strongly affects exoplanet detectability and characterisation, but the presence of exoplanets can also in turn impact stellar RVs!

Finally, this topic is crucial for the preparation of the PLATO mission [launched in 2026 139] and the exoplanet science that will be performed during the mission, since the primary targets are Earth mass planets in the habitable zone around solar type stars (using photometric transits): the RV follow-up of such detections will be complicated by the fact that current techniques do not yet allow to reach very precise mass estimation in such conditions, at least for Earth mass planets. A complete review of processes and tools is given in Watson et al. (in prep).

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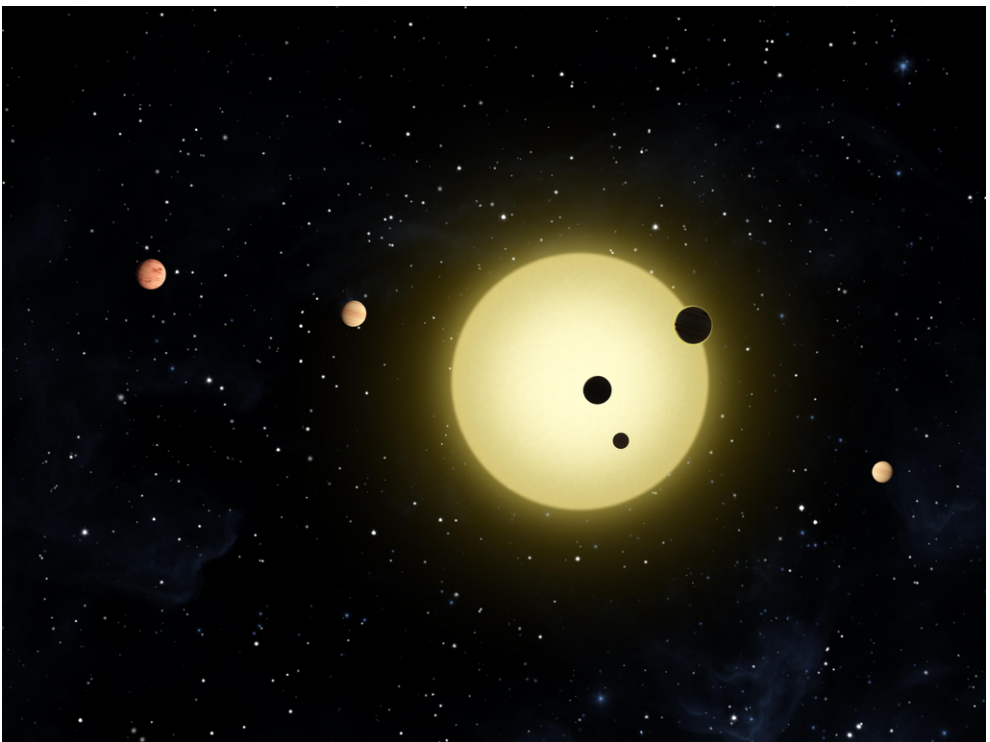
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Stellar activity and transits

by G. Bruno & M. Deleuil



Artist's impression the Kepler-11 system with its transiting planets. Copyright: NASA / Tim Pyle. Licensed under the Creative Commons Attribution-Share Alike 4.0 International. Source: <https://commons.wikimedia.org/wiki/File:Kepler11.png>

Stellar activity and transits

Giovanni Bruno^{1,**} and *Magali Deleuil*²

¹INAF – Catania Astrophysical Observatory, Via Santa Sofia, 78, 95123, Catania, Italy

²Aix-Marseille Université, CNRS, CNES, LAM (Laboratoire d'Astrophysique de Marseille) UMR 7326, 13388, Marseille, France

Abstract. From an observational standpoint, stellar activity poses a critical challenge to exoplanet science, as it inhibits the detection of planets and the precise measurement of their parameters. Radial velocity and transit searches revealed a significant fraction of exoplanet hosts is active, and showed the need to fully understand the different facets of stellar activity and its impact on observables. Moreover, the activity correction is of prime importance for the detection and characterisation of Earth analogues. We present a review of the effects that stellar activity features such as starspots, faculae, and stellar granulation have on photometric and low-resolution spectroscopic observations of exoplanets, and discuss the main aspects of the techniques which were developed to reduce their impact.

1 Introduction

Our understanding of planetary systems is rooted in the knowledge of host stars' parameters and characteristics. Starting from the protoplanetary disc phase, stellar properties affect the types of planets that will form, as well as their composition (e.g. [102, 112]). As systems age, dynamical and magnetic interactions, together with stellar irradiation, are thought to importantly affect the evolution of planetary bodies (e.g. [94]).

The precision obtained on stellar parameters also has a direct impact on the measure of planetary parameters, namely their masses and radii, which are derived relatively to the same parameters of their hosts (e.g. [146]). In turn, the masses and radii of exoplanets are necessary to calculate their mean densities, model their interiors, interpret their transmission spectra and compute atmospheric pressure scale heights (e.g. [34, 53, 67, 153]).

While from one side stellar parameters must be determined with high accuracy and precision, stochastic stellar brightness variations also hamper the precision on measured quantities and prevent a correct determination of planetary parameters and properties. This might result in biases in planetary radii, masses and atmospheric composition.

In this review, the issue of stellar activity from the point of view of photometric and low-resolution transit observations is presented and discussed. In Section 2, we introduce some of the main stellar activity indicators which are used for exoplanet surveys. In Section 3, we show some benchmark cases of active stars observed in the context of exoplanet transit surveys, as well as the main techniques used to constrain the properties of starspots and faculae from observations. Sections 4, 5, and 6 are dedicated to the effect of stellar activity

**e-mail: giovanni.bruno@inaf.it

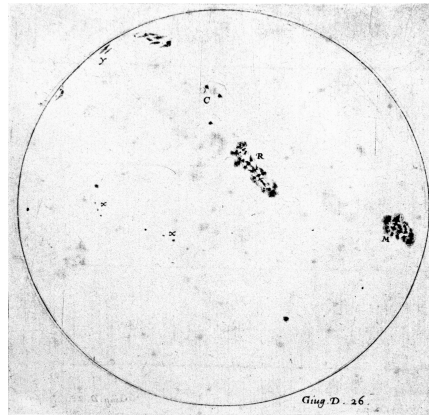


Figure 33: Illustration from Galileo’s *Istoria e dimostrazioni intorno alle macchie solari e loro accidenti* (“History and Demonstrations Concerning Sunspots and Their Properties”, or “Letters on Sunspots”), 1613. Downloaded from Encyclopædia Britannica [128, <https://www.britannica.com/science/sunspot>].

features on transits observed in photometry and low-resolution spectroscopy, and the impact of stellar granulation is the focus of Section 7. Section 8 presents some conclusions and perspectives which can be drawn from what we have learnt so far.

2 Solar and stellar activity observations

The first observations of stellar brightness variations come from the best-studied star we know, the Sun. Hints of the presence of darker regions on the solar disc, which we now call “sunspots”, were recorded in China around 800 BC. In the western world, we can find the oldest records in the drawings in John of Worcester’s *Chronicles*, dating 1128. Thanks to the pioneering use of the telescope for astronomical observations, Galileo Galilei spotted inhomogeneities on the solar disc and recorded them on a plate dating 1612, as shown in Figure 33. The observation of stellar activity phenomena in the visible part of the spectrum, which were later extended to a broader range of wavelengths, enabled also the first detection of a solar flare by R. C. Carrington in 1859.

Pioneering spectroscopic observations of CaII H&K line emission at 393 and 397 nm in solar-type dwarfs were obtained by [124, 180]. Time variations were shown for other spectroscopic lines (e.g. [97]), such as H α . Figure 34 presents CaII H&K lines of a non-active and an active solar-type star, whose spectral line core is affected by strong emission, seen as an indicator of stellar magnetic activity. These activity-related features revealed a dependence of the chromospheric activity level on the spectral type and the rotation period of the star [182]. Main-sequence cool dwarfs have their rotation braked due to the loss of angular momentum through a magnetised stellar wind. Hence, rotational velocity generally decreases with age, as well as activity [63, 143], even if cases were detected of possible weakened magnetic braking and anomalously rapid rotation at old ages [174]. All in all, chromospheric activity is often used as an indicator of age (e.g. [110, 127]).

Photometry also provides indications of time-dependent, stellar brightness variations due to the passage of dark starspots and bright faculae across stellar discs. The amplitude of such variations clearly correlates with spectroscopic indicators (e.g. [9]). Additionally, observations provided evidence that the level of activity decreases with stellar age, corresponding

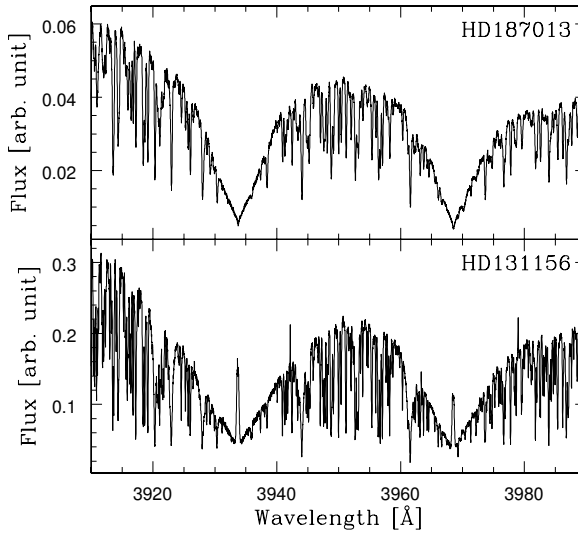


Figure 34: The CaII H and K lines in the SOPHIE spectrum of a non-active (*top*) and active (*bottom*) star. From [29] (their Figure C.1), reproduced with permission ©ESO.

to a transition from spot-dominated to facula-dominated stars [98, 133, 137, 139, 155]. In this regard, our Sun belongs to the category of weakly-active stars [99]. Photometric modulations are observed on various time scales, from days or tens of days – indicating starspots and faculae brought about by stellar rotation – to months or years – revealing the evolution of the brightness dishomogeneities pattern on the stellar disc, activity cycles, and variations in the stellar disc coverage of activity features, or “filling factor” ([119]). This information enables placing constraints on stellar rotation rates, and on the distribution and evolution of activity features, and is therefore particularly helpful to understand the underlying dynamo mechanism across different spectral types (e.g. [25, 26, 30, 35, 68, 71, 90, 154, 165, 166]).

In order to monitor daily photometric variations, ground-based observations are however not sufficient. Day-night alternation forces continuous interruptions, with adverse weather and difficulty to obtain continuous, long-term telescope time worsening the issue. In order to assess the impact of stellar brightness variations on transit observations, long-term uninterrupted monitoring is necessary. Hence, both exoplanet and stellar activity science greatly benefited from the launch of space telescopes dedicated to ultra-high precision and long-duration photometric observations, and better characterised the time scales concerned by stellar activity features (e.g. [11, 30]). Table 2 gives a summary of these space-borne telescopes.

3 High precision photometry

Space missions provided ultra-precise long-duration observations of hundreds of thousands of star, covering nearly all spectral types. However, even for the more than four year-long duration of the *Kepler* observations, the duration of these missions is too short to cover the entirety of a stellar cycle (which, for the Sun, lasts about 11 years). Despite the limited temporal coverage, it was possible to observe a wide variety of brightness variation patterns and to connect them to stellar types. Figure 35 juxtaposes some light curves of the Sun to those of some *Kepler* stars grouped by activity level. Comparative studies of this kind placed the

Telescope	Precision	Time sampling	mag limits	Sample size [stars]	Reference
<i>SOHO/VIRGO</i>	< 10 ppm/min photon noise	60 s	–	The Sun	[61], [60], [79]
<i>MOST</i>	~ 100 ppm/hr	1 – 60 s	$0.4 < V < 6$	5000	[178]
<i>CoRoT</i>	700 ppm/hr	32 or 512 s	$11.5 < V < 16$	170000	[8]
<i>Kepler/K2</i>	80 ppm/hr	59 s or 29 min	$7 < K_p < 17$	~ 170000	[86]
<i>TESS</i>	60 ppm/hr	120 s or 30 min	$4 < V < 12$	> 200000	[140]
<i>CHEOPS</i>	50 ppm/hr	1.05 – 60 s	$6 < V < 13$	> 600	[33]
<i>PLATO</i>	27 ppm/hr (goal)	2.5 or 25 s	$8 < V < 16$	> 10^6	[138]
<i>Ariel</i>	10-100 ppm/hr (goal)	–	bright stars	~ 1000	[169]

Table 2: Space missions dedicated to ultra-high precision stellar photometry and spectrophotometry.

Sun among weakly-active stars, or in a transition phase between active and non-active stars [16, 17, 27, 31, 167]. Moreover, observations on as much as $\sim 10^5$ stars confirmed expectations that cool, mainly convective stars exhibit larger photometric variations (e.g. [64]).

This flux modulations observed in long-term and continuous light curves can reveal surface feature topologies and properties. As an example, Figure 36 shows the light curve of CoRoT-2, hosting one of the first exoplanets discovered by the *CoRoT* mission. From photometric amplitude variations, the size and brightness contrast (i.e., temperature) of starspots and faculae can be extracted. Deviations from a simple sinusoidal shape, which would be an indication of a single activity feature, increase as the number of stellar activity features increases and as their shape gets more complex. We have to remember, however, that we are not able to resolve the stellar disc, so that we cannot distinguish between large stellar activity features and agglomerations of smaller features. In addition, the complex inter-dependence of rotation, differential rotation, and the magnetic field-induced features that evolve in time, produces degenerate cases that hamper the inversion (e.g. [116]). Despite this, long-term data series that allow a time coverage of a few stellar rotations have triggered new developments, bringing new insights into the configuration and evolution of starspot and faculae, but also key constraints on the stellar rotation period and differential rotation.

3.1 Stellar rotation periods

Spots or inhomogeneities on the stellar surface are dragged away by the rotation of the star, resulting in a modulation of its brightness over days or tens of days. This modulation can be used to infer its rotation period, and was taken advantage of by photometric surveys which made large samples analyses possible ([e.g. [1, 16]).

By applying the autocorrelation function (proven to be more robust than the Lomb-Scargle periodogram for this kind of problem), [109] were able to measure the rotation period of more than 34 000 stars, over a sample of more than 133 000 stars. Among other conclusions, they reported an increase of the rotation period with decreasing stellar mass and temperature, broadly consistent with the stellar spin-down law. Their findings, later expanded to an even larger sample (e.g. [95, 145, 149]) also suggest that cooler stars are more evenly covered in brightness inhomogeneities such as starspots – a factor that is particularly crucial to discern the effect of stellar activity on exoplanet transits. This aspect will be better explored in Section 5.

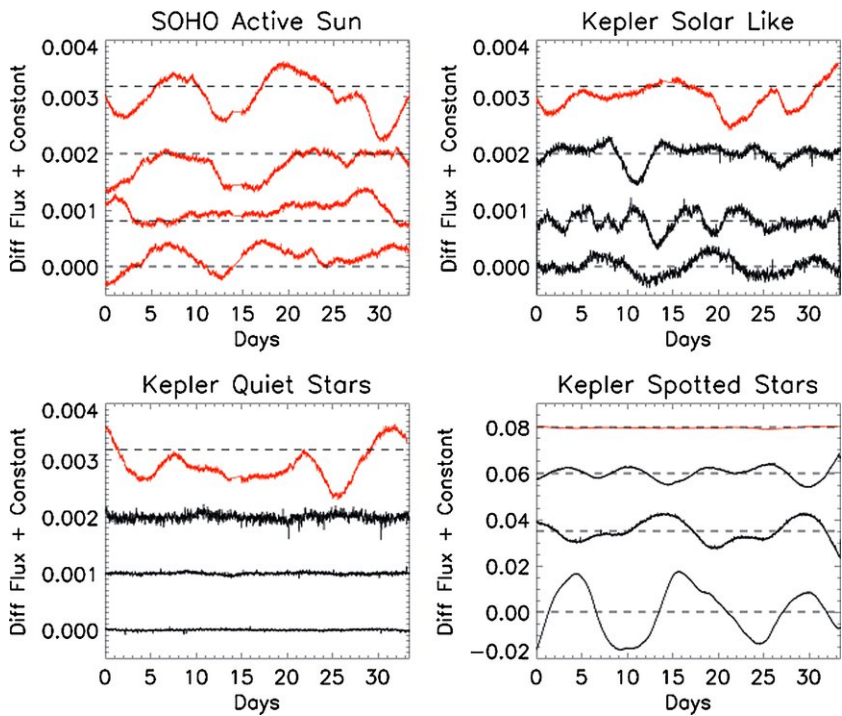


Figure 35: Examples of solar light curves (*upper left*) compared with sample *Kepler* light curves for solar-like (*upper right*), quiet (*lower left*), and spotted (*lower right*) stars. The normalised light curves are offset with respect to each other for clarity. Solar light curves are coloured in red, and *Kepler* light curves in black. In each panel, a solar light curve is shown for comparison. An offset was applied to each light curve. From [16] (their Figure 2). Reproduced by permission of the AAS.

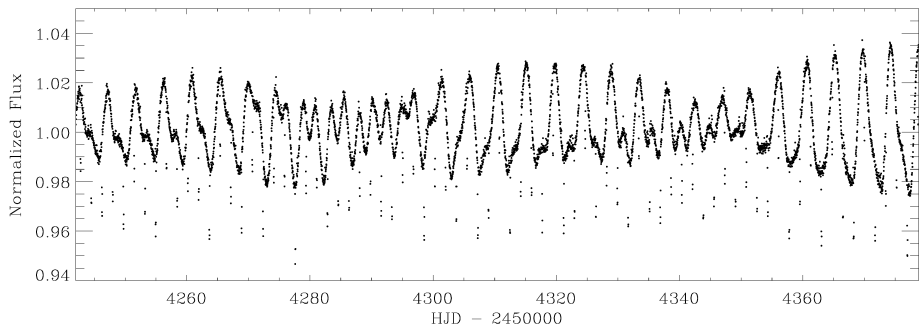


Figure 36: The light curve of CoRoT-2, where the transits of the planet are evident as vertical flux drops. In cases like this, the out-of-transit stellar flux cannot be determined with high confidence. Credit: [6] (their Figure 1), reproduced with permission ©ESO.

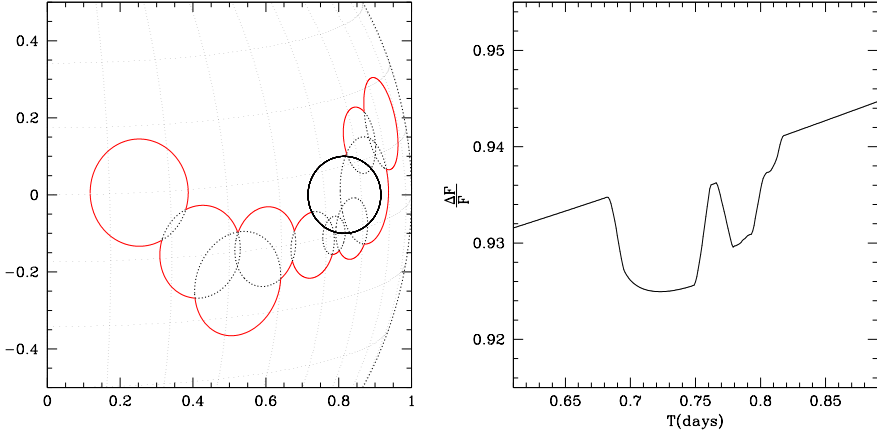


Figure 37: *Left*: Geometric representation of a planet (black) transiting in front of a group of active regions (red). The axes denote the position on the star, normalised to the stellar radius, and are centred on the centre of the stellar disc. *Right*: Corresponding normalised light curve. From [111] (their Figure 2).

3.2 Stellar surface reconstruction

Knowledge on the distribution of stellar surface inhomogeneities can provide very valuable insights on the amount by which transit parameters are affected. Luckily, photometric light curves encode information on the distribution and temperature of such features. The complex problem of light curve inversion has been tackled in a number of ways, since much earlier than it became relevant to exoplanet science (e.g. [38, 56, 141, 175]). Here we report some of the main approaches that have been used in the literature.

- Analytic models describe the stellar disc as the visible part of a sphere emitting a total unit flux, from which the contribution of darker (or brighter) regions with circular shape is subtracted (or added) to calculate the light curve [23, 38, 54, 57, 77, 82, 111]. Usually, brightness inhomogeneities are distorted according to their projection at a given stellar longitude and latitude. Other parameters can be tuned, such as the inclination of the stellar rotation axis, the stellar rotation period, or the stellar limb darkening coefficients. Such a “forward” model is then implemented in a Markov Chain Monte Carlo (MCMC) algorithm (or its variations, such as Nested Sampling) to find the posterior probability distributions for the model parameters (e.g [36, 77, 82]).¹

An analytic model requires only a fraction of a second to be calculated and, in principle, there is no limit to the number of stellar activity features that can be modelled. Clearly, the more complex the model, the longer the computation time required to reach convergence. It is then often desirable to adopt some simplifications, as degenerate cases among parameters might hamper the convergence of the chains. In particular, not all models allow the modelling of planet-starspot occultations, which provide strong constraints on the active region distribution along the transit chord (see Section 4.2). Dedicated modelling approaches were developed for this specific problem [23, 80, 111, 150, 171]. As an example, Figure

¹Among publicly available codes: *macula* (<https://github.com/davidkippling/MACULA>, [82]), *KSint* (<http://edusciisoft.com/KSINT/>, [111]), *Prism* (https://github.com/JTregloanReed/PRISM_GEMC, [171]), and *TOSC* (<http://slnweb.oact.inaf.it/tosc/>, [150]).

37 represents the modelling of a planet occulting a particularly complex combination of bordering active regions.

However, observations provide no constraint on some of the parameters. To overcome this issue, the temperature of the features can be fixed by adopting results from the literature (e.g. [26]), or the stellar rotation axis can be fixed when there is no constraint on such parameter. Moreover, a large number of starspots or faculae might overcomplicate the model, and prevent a proper sampling of the parameter space by the MCMC.

- Alternatively, the stellar visible disc can be modelled with a grid of hundreds of elements, and the contribution of each cell to the total flux can then be evaluated. This allows the modelling of complex shapes for starspots and faculae, but is computationally very expensive (e.g. [28, 125])²

The activity feature distribution can be retrieved in different ways. One is Maximum Entropy Regularisation, where the brightness contrast (or the effective temperature) of starspots and faculae for each cell of the grid is fixed, and a functional that models the light curve given the filling factor f of the cells (i.e., their fraction covered by activity features) is minimised:

$$\Theta = \chi^2(f) - \lambda S(f), \quad (19)$$

where χ^2 is the usual chi-square, S is the entropy functional (a function describing the complexity of the model, so that more complex models are penalised) and λ is a Lagrange multiplier. Despite the large number of free parameters, stable solutions can be achieved with this method, which retrieved active longitudes from some *CoRoT* and *Kepler* light curves of exoplanet-host stars (e.g. [92, 93]). Figure 38 shows the reconstruction of the active region distribution using the light curve of Kepler-17 [92]. Another option is to use a relatively small number of cells to describe the stellar disc: this allows one to adopt variations of the χ^2 statistics, or even MCMC algorithms to achieve the best fit (e.g. [74, 170, 181]).

3.3 Stellar differential rotation and starspot evolution

In order to reconstruct the properties of the stellar dynamo and the generation of magnetic fields in the stellar convection zone, efforts have been directed to collect information on the variation of the stellar rotation period at different stellar latitudes (e.g. [12, 32, 85]). This variation can be expressed as

$$\Omega = \Omega_0 - \Delta\Omega \sin^2 \psi, \quad (20)$$

where ψ is the stellar latitude, Ω the rotation rate at the stellar equator and $\Delta\Omega$ the variation of the rotation rate between the stellar pole and the equator. The “shear” $\alpha = \Omega_0/\Delta\Omega$ is one way to express the deviation of stellar rotation from the one of a rigid body, and is equal to 0.2 for the Sun ($\Delta\Omega_\odot = 0.055 \text{ rad day}^{-1}$).

A variety of methods exist to estimate differential rotation from photometry and spectroscopy (e.g. [26, 91, 120, 144]). The advent of space-borne high precision, long-term photometry added a powerful resource, both for the quantity of continuous data (e.g. [64, 88]), allowing to precisely track stellar rotation periods as a function of stellar latitude, and to asteroseismology [21, 24]. Here we discuss the example of [50], who applied a starspot modelling technique to reconstruct the longitudes and radii of starspots on the *Kepler* light

²In particular, the SOAP 2.0 code [28, 125] allows such a modelling, and is discussed in other proceedings of the EES2019. The code was presented by M. Oshagh at the school (<https://ees2019.sciencesconf.org/resource/page/id/9>) and can be downloaded at <http://www.astro.up.pt/resources/soap-t/> (installation instructions at <https://ees2019.sciencesconf.org/resource/page/id/8>).

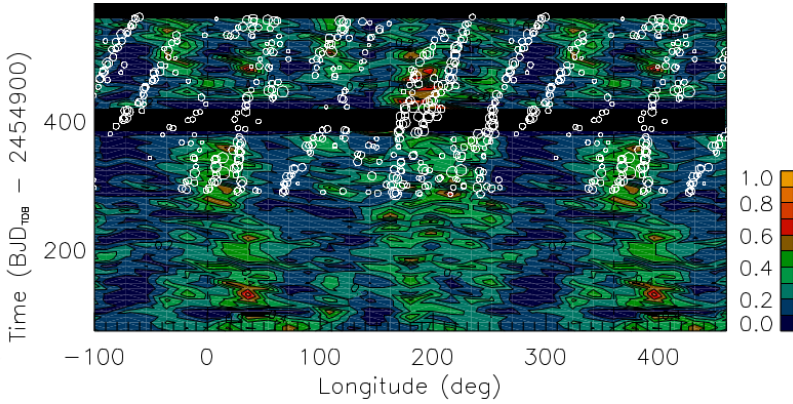


Figure 38: Reconstruction of the active region longitudinal distribution (x -axis) using hundreds of days of Kepler-17 photometric data (time on the y -axis). Modified from [92] (their Figure 9). The colours, from blue to orange, indicate an increasing filling factor, and identify active longitudes. Overplotted in white, the active region distribution identified by [172] using occulted starspots (see Section 4.2). Reproduced with permission ©ESO.

curve of the M dwarf GJ 1243. Using 5 days-long sliding windows and an MCMC sampler, they derived longitudes and radii of two starspot features: a high-latitude feature for over 6 years of observation, and a low-latitude feature evolving in both size and longitude over hundreds of days. For this second feature, they measured a differential rotation rate $\Delta\Omega = 0.012 \pm 0.002 \text{ rad day}^{-1}$ and a shear $\alpha = 0.00114$, one of the lowest values measured for cool stars. The phased light curve map is illustrated in Figure 39.

However, light curve modulations are degenerate with respect to the latitude of stellar activity features, so that these parameters were fixed during the modelling. As discussed in Section 4.2, the occultation of starspots and faculae during transits allows setting constraints on their latitudes, and hence provides significant information for better understanding the dynamo of the host star. Starspot and, for the first time, facula occultations were used to derive the differential rotation rate of a young solar-type star, Kepler-71, by [183]. The result of $\Delta\Omega < 0.005 \text{ rad day}^{-1}$, much lower than that of the Sun's, describes a rigidly rotating star which is set apart from solar-type stars at a similar evolutionary state.

Another major limitation when trying to estimate the configuration of activity features is the description of their evolution, as most of the usual models are stationary. One way to address this issue is to split the light curve in chunks and to model them with different distributions of starspots and faculae [50, 89, 158]. Indeed, stellar activity features do evolve in time, and characteristic laws can be used to describe their growth from first appearance to maximum size, permanence at maximum size and decay until disappearance (e.g. [30, 82, 117, 148, 173]). With a starspot and facula model including feature evolution, longer segments of the light curve can be modelled and the lifetime of individual features (or groups) can be inferred (e.g. [36, 116]).

4 Planetary transits and stellar activity features in photometry

Stellar activity directly affects planetary transits in different ways, as spots and faculae leave their imprint on both their depth and shape. This impacts the observables used to derive the

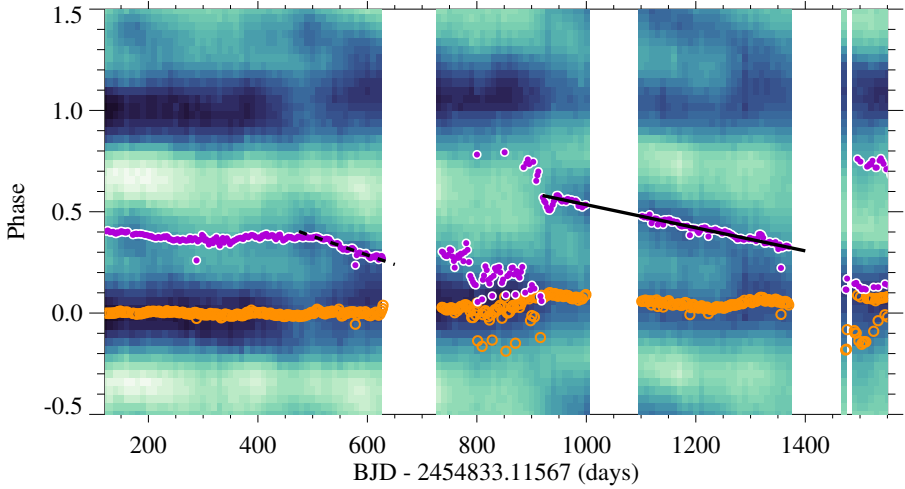


Figure 39: Reconstructed light curve map of GJ 1243, from [50] (their Figure 5). The x -axis corresponds to the observation date, and the y -axis to the stellar rotation phase. The two retrieved starspots are marked in orange and purple. The orange pattern represents the higher-latitude, more stable starspot, and the purple one indicates the lower latitude feature migrating in longitude because of differential rotation. Reproduced by permission of the AAS.

radius of an exoplanet, and as such it is of crucial importance to disentangle them from any kind of contamination.

4.1 Effect of non-occulted activity features on the transit depth

Variations in the apparent transit depth are caused by changes in the out-of-transit stellar brightness. If F_{out} is the out-of-transit stellar flux and F_{in} the in-transit stellar flux (both at the centre of the transit), the transit depth is defined as

$$D \equiv \frac{F_{\text{out}} - F_{\text{in}}}{F_{\text{out}}} = \left(\frac{r_p}{R_\star} \right)^2, \quad (21)$$

where r_p and $R_\star$ are the planetary and the stellar radius, respectively. If the star is covered by activity features, $F_{\text{out}} - F_{\text{in}}$ remains the same, but F_{out} changes with time: in particular, if it decreases (increases) because of starspots (faculae), the transit depth will increase (decrease). Brightness variations are partly accounted for by fitting the out-of-transit flux in proximity of the transit edges with a low-order polynomial, and by computing its value along the transit. The relative effect is more remarkable the larger the brightness modulations are, and the smaller the planet is. Variations can span from a few 10^2 ppm to a few percent [49], with considerable consequences on the inferred planetary properties (such as its internal structure and composition, e.g. [176]). [49] suggested that a different quantity than the transit depth should be adopted in order to minimise the effect of stellar brightness variations. In their notation, $f \equiv F_{\text{out}} - F_{\text{in}}$, $n \equiv F_{\text{out}}$, and the proposed metric is

$$D = \frac{f - n}{p} + 1, \quad (22)$$

where p is the value of the stellar photometric flux unaffected by starspots or faculae. The problem with this expression, however, is that the contribution of faculae to the stellar flux is not known *a priori*, so that the value of p cannot be deduced from observations nor theory.

Another way to estimate the out-of-transit stellar flux is via the use of starspot and faculae modelling. If observations covering a few stellar rotations are available, the distribution of starspots and faculae, as well as their brightness contrast (i.e., effective temperature) can be reconstructed up to some level of confidence, and F_{out} at any intermediate time can be inferred. The precision on this value depends on the uncertainty on the retrieved activity feature distribution and its degenerate solutions. One can remove the transits from the light curve, describe the out-of-transit flux with one of the models outlined in Section 3.2, and use this value for the in-transit flux at any given time. Alternatively, models enabling simultaneous simulations of both activity features and transits can be fitted to the entire light curve. In this way, constraints on both the transit and the activity feature parameters can be obtained, and the need for transit normalisation avoided [36]. The price for this is a considerable increase in computational and model selection effort for the transit fit.

A third option, which has recently gained popularity, is to apply a Gaussian process (GP) regression to the whole light curve altogether, setting the transit as the mean function of the process [3, 14, 59].³ With a GP, the stellar signal is modelled as correlated noise, without the need to describe it explicitly. However, this requires the choice of the functional form for the correlation among observations, i.e. a “kernel”. Some specific kernels are often adopted to model the variations induced by starspots and faculae, such as combinations of the exponential-squared and exponential-sine-squared kernels [69]. The GP can be calibrated so that the kernel parameters (or “hyperparameters”) are used to perform posterior inference on the size, position, and contrast of active regions, as well as their evolution time scales [100, 101]. Convenient open-source software for the GP implementations here described is available online⁴ and can be readily included into an MCMC simulation. GPs are particularly powerful thanks to their flexibility, but they are also very demanding in terms of computation resources: given N data points, this is $O(N^3)$ for most kernels, and $O(N)$ for a specific kind of kernels applied to one-dimensional data sets (e.g. [7, 59]).

Which approach is the most advantageous depends on the case at hand. The simplest model might be the most useful when only a few transits are available (even only one, as it is sometimes the case). Analytic modelling of starspots and faculae is recommended to obtain a physical representation of the starspot configuration and temperature, given an uninterrupted light curve, and it can provide with a more robust assessment of the out-of-transit flux. GPs, on the other hand, avoid the need of precise parametrisation of the geometry of the problem and can be coupled with models of instrumental systematics with a relatively low number of free parameters. None of these methods is completely exhaustive, and the next Section focuses on how occultations of activity features can help break some parameter degeneracy.

4.2 Effect of occulted activity features

When starspots or faculae are occulted by a planet during a transit, they produce typical “bumps” on the transit profile. They affect transit depth measurements in an opposite way to non-occulted ones: if the planet passes in front of a dark starspot, the apparent transit depth will temporarily decrease, and the opposite will happen in case a bright facula is occulted (Figure 40). From the Sun, we know that faculae are often associated to sunspots [81]. They

³A presentation dedicated to GP regression was given by S. Aigrain at the EES2019 (<https://ees2019.sciencesconf.org/resource/page/id/9>).

⁴For example, *George* (<https://github.com/dfm/george> [7]), *celerite* (<https://github.com/dfm/celerite>, [59]) and the GP regressors in *scikit-learn* (<https://scikit-learn.org>, [129]).

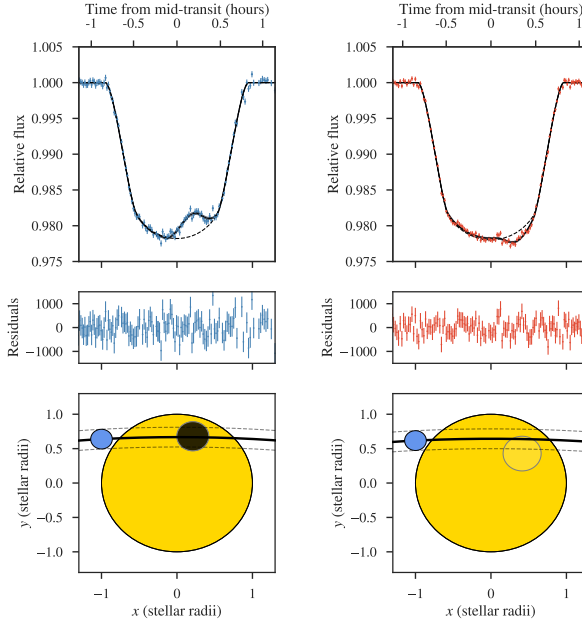


Figure 40: Effect of the planet crossing of a spot (*left*) and a facula (*right*) on the transit profile. From [58](their Figure 7).

are mostly visible when they are close to the stellar limb, and are therefore described by a wavelength-dependent limb brightening law (e.g. [156, 164]).

If the temperature contrast of the activity features with the average stellar photosphere is large enough, given the noise level of the data, their occultations can be easily identified in the transit profile. In principle, they can therefore be removed from the transit profile, which can then be fitted with a standard transit model. However, if their signal cannot be clearly singled out, or during a given single transit the planet crosses multiple stellar activity features, removing the imprint of the occultation might be problematic. The case of undetected occulted features was discussed by e.g. [10], while the problems of multiple starspot crossings in the same transit were explored by e.g. [49] and [158] on the benchmark case of the hot Jupiter CoRoT-2b. These latter authors found that, under the assumption that all crossed features are dark starspots, the transit radius of CoRoT-2b might be overestimated by up to 3% if the crossing effect is not taken into account. However, our knowledge of starspots and faculae in other stars is still not complete enough to exclude the contribution of faculae, which could average out this effect, or even increase the apparent transit depth in individual transits. Indeed, [36] found indications that some of the deepest transits of CoRoT-2b, previously considered as the closest ones to the “unperturbed” transit profile, might be affected by faculae.

A very interesting advantage of starspot and facula crossings is that their position on the stellar disc (assuming the stellar rotation axis is known) can be determined with a much larger precision than in the case of non-occulted features. Provided a stellar inclination angle of 90° with respect to the line of sight, the feature latitude ϕ can in fact be constrained within the

stellar latitudes crossed by the transit chord:

$$\phi = \arctan\left(\frac{r_p}{R_\star}\right). \quad (23)$$

Thanks to this, the latitudinal distribution of the starspots on HAT-P-11 was shown to follow a solar-like behaviour [115].

Occultations also provide precise constraints on the longitude of active regions, which are found to be consistent with the reconstruction of the stellar surface using only unocculted features (see Figure [92]). In the special case of Kepler-7b, [52] were able to observe the recurrent passage of the same group of occulted starspots on the transit profile. Together with a previous measurement of the stellar obliquity via the Rossiter-McLaughlin effect, and helped by the commensurability of the stellar rotation and planet revolution period ($P_\star/P_{\text{planet}} \simeq 8$), it was possible to follow the starspots moving across different transits. Thanks to the long duration of the Kepler-7 light curve, starspot groups were observed without interruption for about 100 days, and important constraints on the lifetime of these features were obtained.

A last aspect to underline is the possibility of constraining the size and temperature of spots and faculae thanks to their occultation. A minimum value on their size can be inferred from the duration of the bump they produce on the transit profile, and this can be scaled to a feature size if the transit duration and stellar radius are known. With a simple geometric model that takes the activity feature brightness contrast with the stellar surface as a free parameter, one can derive the temperature contrast of the active feature to the stellar photosphere by inverting the ratio between black body laws [157]:

$$\rho = \frac{\exp(h\nu/K_B T_{\text{eff}}) - 1}{\exp(h\nu/K_B T_{\text{feat}}) - 1}, \quad (24)$$

where ρ is the brightness contrast with the stellar photosphere, h is Planck's constant, K_B is Boltzmann's constant, ν is the frequency of observation and T_{eff} and T_{feat} are the stellar and the activity feature effective temperature, respectively. This method was adopted in several cases (e.g. [105, 158]).

Clearly, the black body law is only a useful approximation as stars, starspots, and faculae are not black bodies at different temperatures, and each one has its own spectrum and specific absorption lines. Section 5 contains a discussion on how transmission spectroscopy can help us probe not only planetary spectra, but also those of occulted stellar activity features.

5 Starspots and faculae in transmission spectroscopy

Transiting exoplanets are key targets to probe the physical structures and chemical compositions of their atmospheres. As their upper atmosphere absorbs and scatters starlight, transmission spectroscopy can be used to infer the physical conditions within their atmospheres by observing the transit depth at different wavelengths [43]. Here, also, the impact of starspots and faculae cannot be neglected because, as discussed in Section 4.1, they increase and decrease the measured transit depth, respectively. The variations are not only time-dependent, but also wavelength-dependent: according to Wien's law, the peak emission from a black body moves to longer wavelengths as the black body gets cooler, and there is more sensitivity close to the emission peak. Hence, the effect of starspots and faculae is stronger in the visible, and less prominent (but still important) in the infrared (IR).

As shown in the example of Figure 41, it was found that unocculted starspots (faculae) result in a rising (descending) slope in the transit depth from the near-IR to the ultraviolet, which could be wrongly interpreted as Rayleigh-scattering dust (or its absence) in the planet's

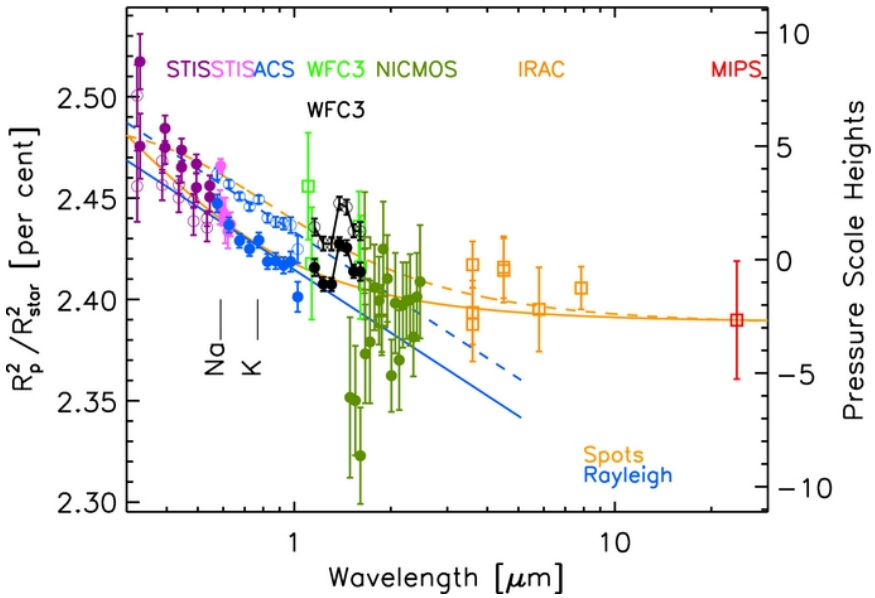


Figure 41: The transmission spectrum of HD 189733 b in different bands. In the visible part of the spectrum, blue and orange colours represent Rayleigh scattering and unocculted starspot models, respectively. From [108] (their Figure 4). ©AAS. Reproduced with permission.

upper atmosphere (e.g. [4, 108, 130, 132, 134, 162]). The opposite happens for occulted active regions [126].

Among molecular spectral signatures which are nearly routinely detected, the water absorption band is the main one observed by *Hubble Space Telescope*'s Wide Field Camera 3 instrument, and is the main indicator of oxygen abundance and metallicity in an exoplanet atmosphere. However, in starspots cooler than ~ 3000 K, water vapor could exist [179] and produce absorption features that can mimic planetary water absorption at ~ 1.4 and $\sim 2.3 \mu\text{m}$ (e.g. [177]).

Because of this, an assessment of the activity level of the exoplanet host is necessary before any interpretation of a transmission spectrum is attempted. The correction of the activity-induced transit depth variations can be done by a photometric monitoring of the host star, with stellar models (such as [39] 1D ATLAS and PHONENIX [76]) to account for the wavelength-dependent flux correction, in order to properly evaluate the stellar contribution (e.g. [4, 75]). In this kind of modelling, the size of the activity features was shown to be essential [135]. With the use of stellar synthetic spectra to model both stars with different spectral types and their starspots, [135] found that active regions with angular size comparable to the size of large sunspot groups ($\sim 2^\circ$, [106]) can produce up to tens of percent variations in atomic and molecular absorption bands, against a few percent in the case of features which are four times as large. This is due to the fact that, for a given activity-induced stellar photometric variability amplitude, small active regions require much larger covering fractions to produce such amplitude, and produce a larger contamination in the transmission spectrum. Figure 42 presents their findings on spectral contamination from starspots on M dwarfs. With the use of similar simulations, the same effects were shown to be much smaller for FGK stars [136].

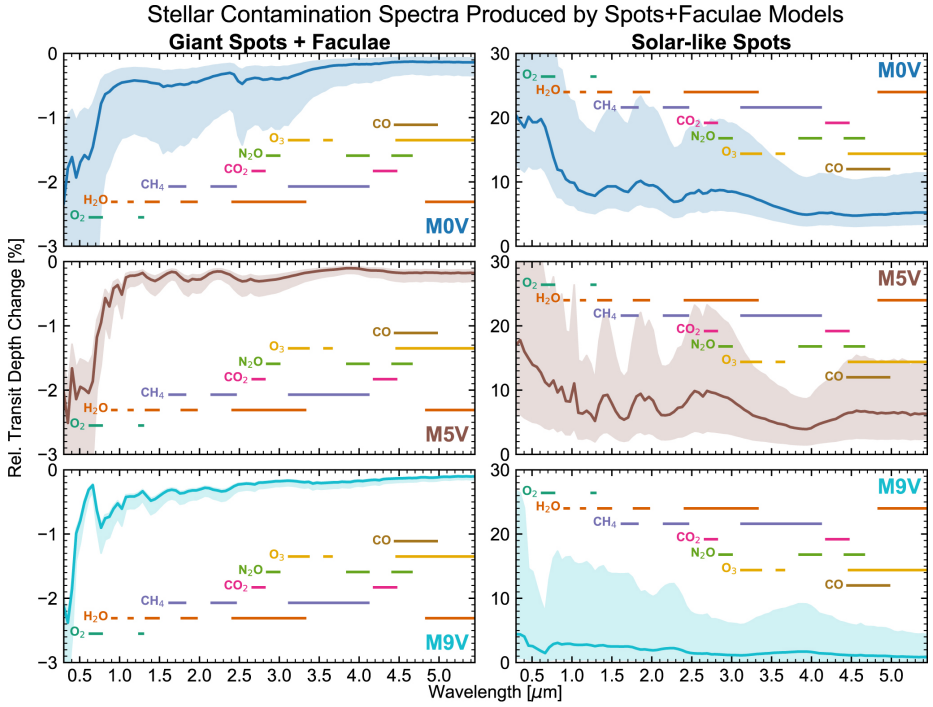


Figure 42: The amount of contamination of transmission spectra for planets hosted by active M stars, in the case of giant (*left*) and solar-size starspots (*right*), according to [135] (their Figure 7). The molecular absorption bands at various wavelengths are reported. Reproduced by permission of the AAS.

The shape of the out-of-transit light curve could be the key to distinguish giant from solar-like activity features on planet hosts, and so to identify optimal targets for transmission spectroscopy of planets orbiting cool stars. In the case of giant starspots, wide variations of the light curve will be observed, with a periodicity which depends on the stellar rotation period. In the case of small starspots, especially if these are uniformly distributed on the stellar disc, brightness variations will have a much smaller amplitude and no clear periodicity. This might be the most frequent case for M stars (see Section 3) where, on average, a small starspot going out of view could be compensated by another one coming into view. The two scenarios could be discerned thanks to a continuous monitoring of the host star in the time windows during which transits are observed. Inspection of the light curve for at least two or three rotation periods in different spectral bands should enable to distinguish the type of light curve variations, and so to gauge the amount of contamination on the transmission spectrum. Remarkably, a time-evolving, longitudinal map and effective temperature reconstruction of the active regions on WASP-52 was obtained from ~ 600 days of *BVR*I multiband photometry, implying a reduction of their effect on transmission spectroscopy by about one order of magnitude [142]. A complementary approach would be to model the temperature and size of active regions using out-of-transit low-resolution spectroscopic observations. This technique showed promising results for the correction of transmission spectra affected by large-contrast spots, in modelled bright ($V = 9$) stars [47].

The time dependence of the “filling factor” (i.e., the fraction of the stellar surface covered by activity features) also operates a variation in the baseline between two or more transit observations. The photometric monitoring of the host star’s activity level was, in fact, shown to be key to correctly stitch together multi-epoch transmission spectroscopy observations, aiming either at extending the wavelength coverage or at detecting temporal variations of the planet’s atmosphere. Retrieval exercises including the contribution of activity features to the spectrum were also shown to provide constraints on both the atmospheric and the starspot parameters (e.g. [15, 37, 134, 184]).

Simultaneous transit observations in multiple wavelengths can provide very useful information on the physical properties of activity features when they are occulted by the planet. In that case, indeed, the relative flux increase (the height of the starspot occultation bump) Δf measured in different wavelengths λ probes specific absorption lines. This makes it possible to place constraints on the effective temperature of the activity feature, whose emission is modelled using stellar synthetic spectra. Following [160],

$$\frac{\Delta f(\lambda)}{\Delta f(\lambda_0)} = \frac{1 - F_{\lambda}^{T_{\text{feat}}} / F_{\lambda}^{T_{\text{eff}}}}{1 - F_{\lambda_0}^{T_{\text{feat}}} / F_{\lambda_0}^{T_{\text{eff}}}}, \quad (25)$$

where F_x^y is the normalised flux measured at wavelength x and effective temperature y , and λ_0 is a reference wavelength. Figure 43 shows the so-determined spectrum of a starspot occulted during a transit of HD 189733b. For the specific case of this active K0-dwarf star, [160] derived a spot temperature $T_{\text{spot}} \approx 4250 \pm 250$ K and spot-to-star temperature differences compatible with what is measured for sunspots. In the case of WASP-19, the fit of the occultation distortion at varying wavelengths allowed the determination of better than 100 K constraints on a bright and a dark active region [58].

6 Effects on other transit parameters

Among the transit parameters which are affected by stellar activity features, limb darkening (LD) is one of the most crucial. In turn, centre-to-limb stellar brightness variations modify both the transit profile and its apparent depth, affecting the determination of the exact radius of the planet [48, 70, 107]. In Figure 44, left panel, the transit profile shape variation is shown when the effect of LD is included. In the figure, the transit depth increase is dubbed “overshoot”. On the right, the overshoot is analytically derived and shown for different spectral types. If stellar activity features affect the LD coefficients, the transit depth measure (and therefore the measured radius of the planet) for a given stellar type is prone to error.

Usually, LD coefficients are obtained by fitting stellar centre-to-limb specific intensity model profiles with a variety of increasingly complex analytic relationships. For example, two popular laws are the quadratic [87] and the non-linear [45] LD law. Using θ as the angle between the stellar surface normal and the line of sight, they are parametrised as

$$\frac{I_{\lambda}(\mu)}{I_{\lambda}(1)} = 1 - a_1(1 - \mu) - a_2(1 - \mu)^2, \quad (26)$$

and

$$\frac{I_{\lambda}(\mu)}{I_{\lambda}(1)} = 1 - \sum_{n=1}^4 a_n(1 - \mu^{n/2}), \quad (27)$$

respectively, where $\mu = \cos \theta$ and $I_{\lambda}(\mu)/I_{\lambda}(1)$ expresses the ratio between the specific intensity at any wavelength λ and position μ and the centre of the stellar disc (corresponding to $\mu = 1$).

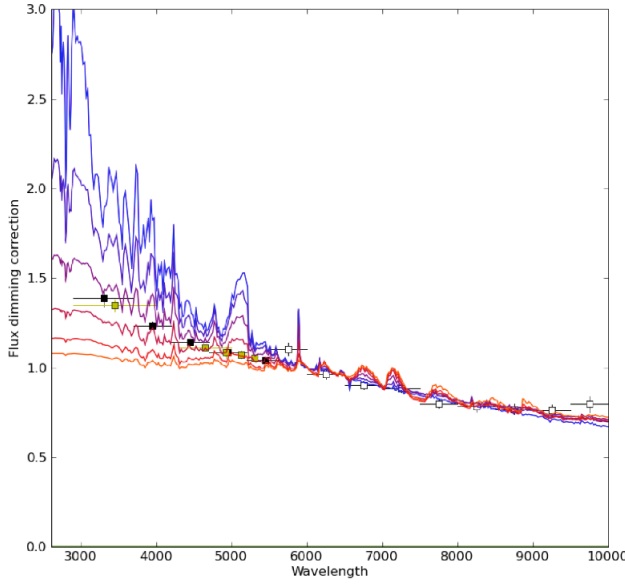


Figure 43: Relative height of the spot occultation bump with respect to the same quantity at 600 nm, as a function of wavelength (data points), and Kurucz 1D ATLAS models used to fit the measured flux rise during a transit of HD 189733b. Measurements (symbols) were obtained from *Hubble Space Telescope*’s STIS G430L and ACS instruments. Starspot models range from 4750 to 3500 K in 250 K intervals (blue to orange, respectively), while a $T_{\text{eff}} = 5000$ K ATLAS stellar model [39] was used for the host star. From [160], MNRAS 416, 1443–1455 (their Figure 10).

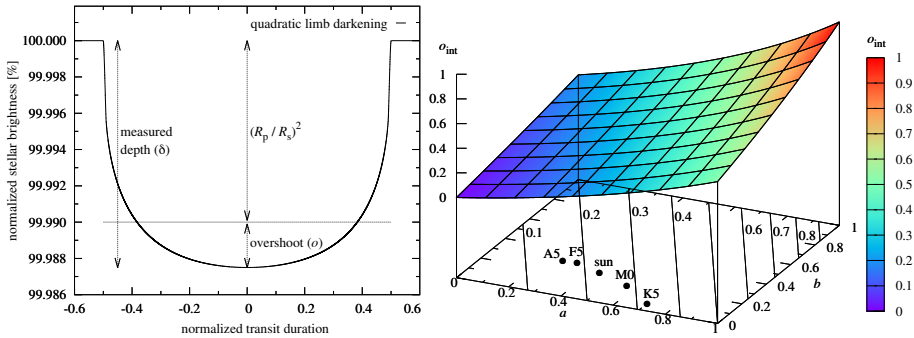


Figure 44: *Left*: the effect of stellar limb darkening on the transit shape and depth. The increase in transit depth over the planet-to-star surface ratio is here called “overshoot” (o). *Right*: The variation in overshoot (o_{int} , z -axis) for a quadratic limb darkening law (coefficients a and b) and orbit inclination $i = \pi/2$ rad for different stellar types. Variations in the limb darkening coefficients due to stellar activity cause variations in the overshoot. From [70] (Figures 1 and 3), A&A 623, A137, pp. 2 and 5, reproduced with permission ©ESO.

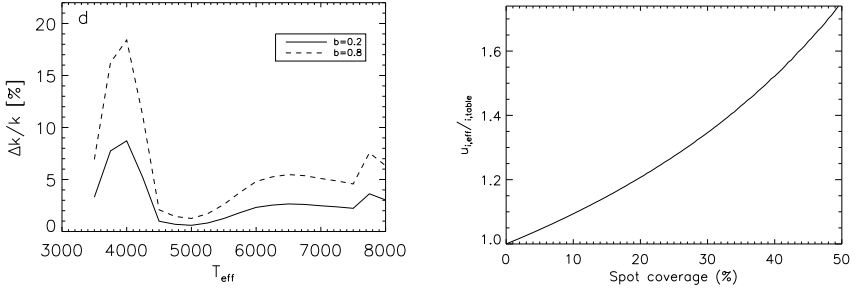


Figure 45: *Left:* Variation of $k \equiv r_p/R_\star$ from the overshoot given by [159] and the [46] LD coefficients, for a low (continuous line) and grazing (dashed line) impact parameter b . *Right:* variations in the observed LD coefficients for varying starspot filling factors for a star with $T_{\text{eff}} = 5775$ K and $T_{\text{spot}} = 3775$ K. The positions of the spots were randomly chosen on the visible hemisphere. Figures from [48], A&A 549, A9, pp. 4 and 8 (their Figures 2 and 5), reproduced with permission ©ESO.

The LD coefficients a_n can be fitted to the stellar models and then used as fixed parameters to model the transit profile [45]. In particular, the non-linear law has been shown to produce the most accurate fit to the models so far (e.g. [113], and references therein).

A fit of the LD coefficients to the transit profile is often more desirable than fixing them to theoretical values. As illustrated in Figure 45, left panel, this is due to the unavoidable uncertainties in the stellar parameters, as well as to differences in the stellar models that are used. This is particularly the case when stellar activity features produce time-dependent slight deviations of the stellar parameters, that affect the LD coefficients and therefore the transit depth overshoot. When the LD coefficients are included in the transit fit as free parameters, however, the non-linear law often hampers convergence of the MCMC sampling of the transit parameters, and simpler laws need to be used. Starspots were found to potentially produce a few percent variations in the quadratic LD coefficients, as well as up to $\approx 10\%$ variations in $r_p/R_\star$, depending on the stellar effective temperature and transit impact parameter [48]. The maximum effect was found around $T_{\text{eff}} = 4000$ K, grazing transits, and for large spot filling factors, as shown on the right panel in Figure 45. Flux-transport models based on the Sun showed that the effect of faculae can also be non-negligible, and that deriving LD curves from the transit fits with minimum residuals induces wavelength-dependent biases due to both spots and faculae [151].

Because of the limited accuracy of the quadratic LD law, some variations were proposed, as well as efficient calculation methods and proper priors on the coefficients to include in the transit fit (e.g. *kippling2013, espinoza2016, morello2017, claret2018, maxted2018*). More accurate three-dimensional stellar models also help us better understand the impact of granulation and the behaviour of LD with fundamental parameters (e.g. [103]). Due to variations on individual stars and the time dependence of the activity pattern on the stellar photosphere, however, analysing exoplanet transits still requires a compromise between accuracy and computation efficiency when adopting an LD law to implement in an MCMC.

As previously highlighted, LD is a wavelength-dependent effect: it is stronger in the visible, but its effects are also measured in the IR. It has therefore to be taken into account both in transit photometry and spectrophotometry. Observations with the *Hubble Space Telescope* deserve a particular mention: objects which do not lie in the continuous viewing zone of *HST*

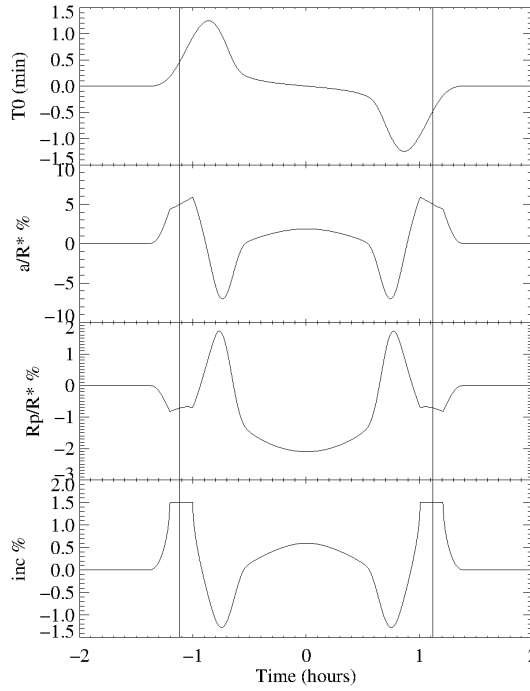


Figure 46: Simulated difference between a transit with and without starspot occultation, as a function of the mid-starspot time compared to mid-transit time, for the parameters of WASP-10b. From top to bottom: transit mid-time, orbital semi-major axis divided by stellar radius, planetary-to-stellar radius ratio and orbit inclination. From [13] (their Figure 10).

cannot generally be observed for a full continuous transit. Sometimes, the transit edges are sacrificed in order to obtain a precise measure of the transit depth. As a consequence, LD coefficients cannot be fitted, but are fixed to tabulated values obtained from stellar models at different wavelengths and in the appropriate filters (e.g. [177], and references therein). The *James Webb Space Telescope* will instead be placed at the Sun-Earth L2 point, and will not suffer from pointing interruptions. For this observatory, the wavelength dependence of the LD coefficients will be fully appreciated.

Stellar activity features have been found to influence the measure of other transit parameters, too. A notable case is the one stellar density, which can be constrained to a high level of precision from planetary transits [131, 163]. Studying the light curve of CoRoT-7b, where the individual transits are not fully resolved, [96] discussed an underestimation of the stellar density from the transit fit compared to the derived spectroscopic value, likely due to stellar variability. Other studies [5, 13, 125] found effects on the apparent transit timing. Figure 46 shows the simulated difference between a transit with and without starspot occultation, as a function of mid-spot time compared to mid-transit time, for the hot Jupiter WASP-10b [13]. The largest effect is produced when the starspot bump is close to the transit edges, marked with vertical lines. In particular, variations in the mid-transit time (top panel) across multiple transits may mimic transit timing variations (TTVs), which might be misinterpreted as due to non-transiting companions [2]. [13] found hints of starspot contamination by observing a large number of transits with significant variations in the depth of individual transits.

This latter is another indication that the study of planets around active host stars deserves a special attention. It requires a careful monitoring of the star to evaluate and correct for the contribution of the starspots to the stellar flux due to stellar rotation and their changing characteristics and location.

7 Planetary transits and stellar granulation

Stellar brightness inhomogeneities caused by stellar convection are another source of noise for transit parameter determination. Granulation produces variations on time scales from tens of minutes to hours. Hot plasma is uplifted towards the stellar surface and produces brighter areas, called granules by [51] for the first time. Cooler plasma sinks between granules, and produces separating lanes between granules [122, 123]. The first observations on granulation were carried out on the Sun [72], but this phenomenon is now thought to be common among Sun-like stars (e.g. [40, 55, 66, 121, 152], and references therein).

7.1 Effect on measured stellar parameters

The first direct effect of stellar granulation is on the stellar parameters, which are fundamental in order to extract the planetary radius from transit observables. The power of the granulation signal is correlated with the stellar surface gravity g : as a star evolves from dwarf toward the giant phase, its g decreases; at the same time, its convection zones deepen and the granulation timescale increases. By measuring the photometric variations on a time scale of 8 hours, [18] highlighted brightness variations driven by granulation. The stars of this study span an effective temperature between 4500 to 6750 K, $\log g$ between 2.5 and 4.5 (cgs units), and relative brightness variations lower than 10^{-3} . Taking advantage of g measurements obtained from *Kepler* asteroseismology observations [42], they derived the “flicker sequence” of stellar evolution. This sequence shows a tight correlation between the granulation-induced brightness variation and the stellar $\log g$ (see Figure 1 of [18]). Within this phase space, [18] fitted an empirical relation (extended by [19]) to infer g to within 25% from the granulation flicker of inactive Sun-like stars, from their main-sequence to their giant phase of evolution.

In addition to surface gravity for the determination of stellar radii, stellar density is also particularly useful to correctly derive transit parameters. This could be carried out via the so-called “asterodensity profiling” [84, 161]. By using 588 catalogued *Kepler* target stars with asteroseismology measurements [41, 73], [83] reported a linear trend between the stellar density and the 8 hr stellar flicker for stars with $4500 \text{ K} < T_{\text{eff}} < 6500 \text{ K}$, $3.25 < \log g < 4.43$, and K_p magnitude ≤ 14 . The fitted relation has a model error in the stellar density of about 32%: this is an $\sim 8\times$ lower precision than asteroseismology, but the relationship can be applied to a $\sim 40\times$ wider sample of targets with no asteroseismic measurements. Capitalising on this, [20] re-estimated the stellar radii of 289 bright ($K_p < 13$) candidate planet host stars with $4500 \text{ K} < T_{\text{eff}} < 6650 \text{ K}$ and assessed the impact on the parameters of their planets. They found the $\log g$ derived from broadband photometry and spectroscopy to be systematically lower than the one obtained from the flicker-based calibration (~ 0.2 dex median difference, with RMS = 0.3 for spectroscopy and 0.4 for broadband photometry). Hence, nearly 50 of the brightest stars in their sample resulted to be subgiants and their radii, together with those of their planets, on average 20%-30% larger than previously measured.

7.2 Effect on transit parameters

Stellar granulation affects the measured transit depth, duration determination of transit ingress and egress, and LD parameters. To quantify the effects, [44] used three-dimensional radiative-

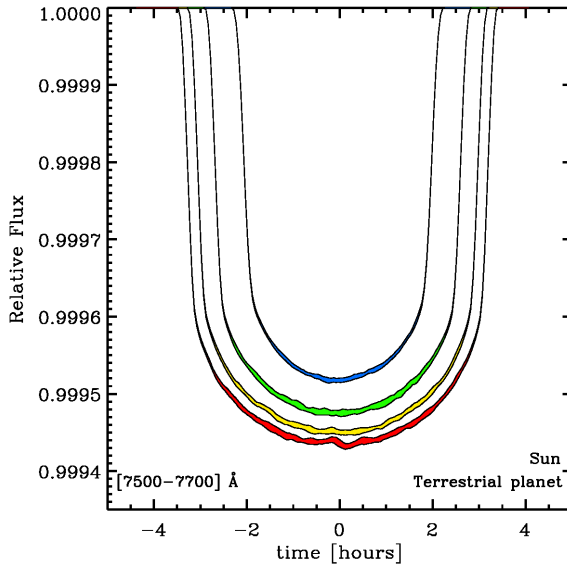


Figure 47: Simulated transit profile variations of a terrestrial planet transiting the Sun, due to different realisations of the modelled granulation pattern. Filled contours indicate the deepest and the dimmest transit profile for different orbital inclinations i in the 750-770 nm band. Blue corresponds to $i = 90.85^\circ$, green to $i = 90.65^\circ$, yellow to $i = 90.45^\circ$, and red to $i = 90.25^\circ$. Credit: [44] (their Figure 11), reproduced with permission ©ESO.

hydrodynamical stellar models from the *STAGGER* grid [104] to model the transit of a hot Jupiter, a hot Neptune and a terrestrial planet in front of the Sun and of a K star.

The simulations were run by considering several convective turnovers and by taking into account temporal variations in the granulation intensity (e.g. about 10 minutes for the Sun, [118]). They were also performed in several wavelength windows, covering the bandpasses of most instruments used to observe spectrophotometric transits. The modelled root mean square of the stellar flux in the visible (between 1 and 16 ppm) was close to the observed photometric variability of the SOHO quiet Sun data (10 to 50 ppm: [62, 78]) and was the largest in the visible. Then, different realisations of the stellar irradiance calculated from the granulation patterns were averaged in order to derive the photometric noise they produced. This resulted in detectable variations of $r_p/R_\star$, larger for the Sun than for the K star. In the visible, for example, they observed up to 0.90% and up to 0.45% uncertainties in the planet radius for the Sun and the K star simulations, respectively. Also, larger uncertainties were found for terrestrial than for giant planets. Figure 47 illustrates the case of a terrestrial planet transiting in front of the Sun, in visible wavelengths, for different orbital inclinations. Given the required precision on planetary radii of a few percent at most, in order to be able to significantly constrain planetary core sizes [176], such estimates imply that granulation must be considered a non-negligible source of uncertainty. [168] quantified the effect of granulation on the transit parameters, which can induce errors of up to 10% on the ratio between the planetary and stellar radius ($r_p/R_\star$) for an Earth-sized planet orbiting a Sun-like star. Stellar flicker is already included in the error budget for the *CHEOPS* mission [33], and

it will be even more so for the *PLATO* mission [138], which is designed to achieve a 2% precision on planetary radii.

The apparent radius variation was also shown to be dependent on the orbital inclination of the planet, with additional effects on the duration of transit ingress and egress and correlations with the LD parameters. [114] explored the correlations among these effects and found uncertainties on the planetary radius up to 3.6% in the *PLATO* band, due to granulation and stellar oscillations only. These authors highlighted the importance of independent constraints on the transit impact parameter, or of follow-up observations at longer wavelengths, where LD is weaker, to reduce the significance of this effect. On the other hand, the method used to retrieve the transit parameters seems crucial in the determination of the error budget: Gaussian processes, which are particularly suitable for the modelling of a stochastic phenomenon such as stellar granulation, showed to provide encouraging results with specific kernel choices [14].

In the context of the forthcoming *James Webb Space Telescope* (e.g. [22, 65]) and the future *Ariel* mission [169], the potential impact of stellar granulation and oscillations on transit spectrophotometry has started to be evaluated. [147] reported that the expected effect is weak, especially in the near-IR. There, its impact could even be negligible, but with a dependency on the atmospheric scale height and the transit duration. As a case which is most likely to be impacted by granulation, these authors mention the one of ‘a terrestrial planet with a secondary atmosphere orbiting a nearby Sun-like star on a long period’ (e.g. [147], p.2877).

8 Conclusions

In this review, we presented the main issues caused by stellar activity on exoplanet transit observations, as well as the main approaches which are used to correct for them. In several cases, the combination of ultra-precise measurements and theoretical modelling is required to precisely quantify the possible biases on planetary parameters. Depending on the observations and the scientific goal, complementary observations can provide a useful way to better assess the effects of stellar activity. On the other hand, attention to the determination of the physical properties of stellar activity features such as starspots, faculae and granulation and of their temporal evolution has grown in the past few years. The goal of a few percent precision in planet radii, and few tens of parts per million in transmission spectroscopy, is challenged by the imprints of stellar activity on the observables, and requires these phenomena to be taken into account as an unavoidable component of star-planet observations.

As we push our search towards terrestrial planets, the understanding of stellar activity and its proper correction becomes an increasingly necessary complement to improved instrument performance. The best outcomes will then result from synergies between different observation techniques and modelling perspectives. Thanks to these, we will be able to obtain as much precise as possible insights into the physical properties of the most interesting discovered systems. At the same time, this race towards small size planets with precise and accurate parameters will help refine our knowledge and understanding of stars.

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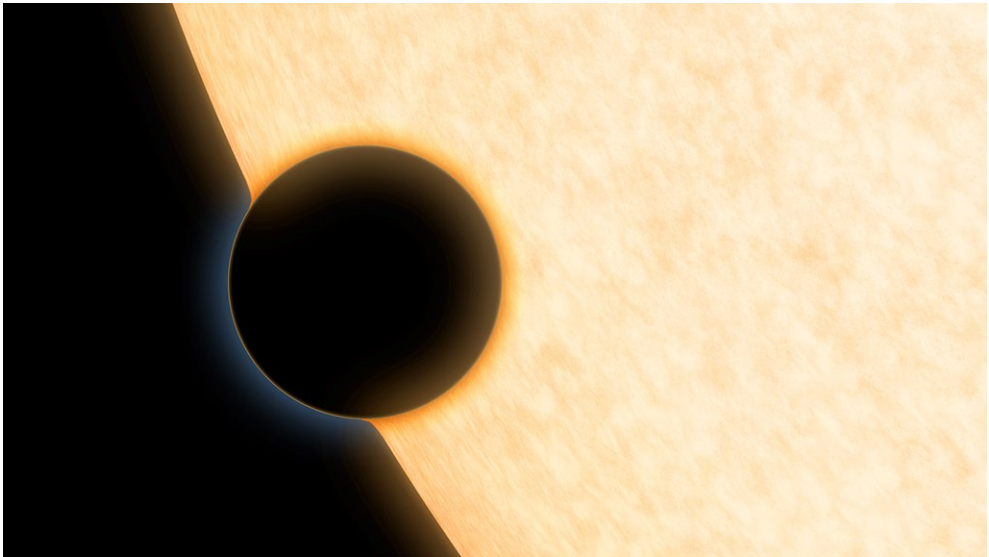
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High resolution spectroscopy for exoplanet characterisation

by M. Brogi



Artist's impression of the stellar light passing through the transit planet atmosphere. Copyright: NASA/JPL-Caltech. Licensed under the Creative Commons Attribution-Share Alike 4.0 International. Source: https://commons.wikimedia.org/wiki/File:Artist_Illustration_of_planet_HAT-P-11b.jpg

High resolution spectroscopy for exoplanet characterisation

Matteo Brogi^{1,†}

¹Department of Physics University of Warwick, Coventry, CV47AL, UK

Abstract. The last two decades have seen an unprecedented rise in the number of known planets around other stars (exoplanets), and the establishment of observational techniques to probe the chemical and physical properties of their atmospheres. This Chapter focuses on one of these techniques, high-resolution spectroscopy with ground-based telescopes. It covers the key aspects of the method and how it is different from other standard techniques, the science achieved so far, current challenges both on the observational and modelling side, and the prospects for future observations of smaller and cooler planets.

1 Introductory concepts

1.1 The underlying scientific questions

Two decades of discoveries have delivered over 4,000 confirmed exoplanets. Given this large population, it is tempting to start comparing their characteristics with those of the planets we see in the solar system. However, the comparison is far from straightforward. Looking directly at the planets *detected* the situation seems unfavourable for us, as the parameter space occupied by our solar system is relatively empty (Figure 48). A prominent class of planets emerges instead, which groups gas giants with approximately the mass of Jupiter but orbits well within the orbit of Mercury. These are called *hot Jupiters*, and were the first exoplanets discovered. As odd as these planets may seem, they are particularly welcome by the community as they are not only easy to find, but also to study, as we will directly estimate in Sections 1.5.1–1.5.4. Hot Jupiters also have a simplified interpretation, as their high temperatures quickly drive the chemistry towards equilibrium, and the absence of major cold traps allows us to relate any species we recognise in the spectrum to the actual composition of the atmosphere, unlike the solar-system giants where the signatures of important molecules such as water vapour remain buried below the observable level of the planet’s atmosphere.

However, the observed exoplanet population cannot just be considered at face value, as detecting smaller planets further away from their parent stars becomes increasingly challenging. A careful exercise must be performed, accounting for the observational biases, to reveal the true underlying exoplanet population. When doing so, we find out that hot Jupiters are actually rare, and that the most common exoplanets have instead size between those of the Earth and Neptune [30, 63]. Once again, this strikes us as odd as there aren’t any planets like those in our solar system.

[†]e-mail: m.brogi@warwick.ac.uk

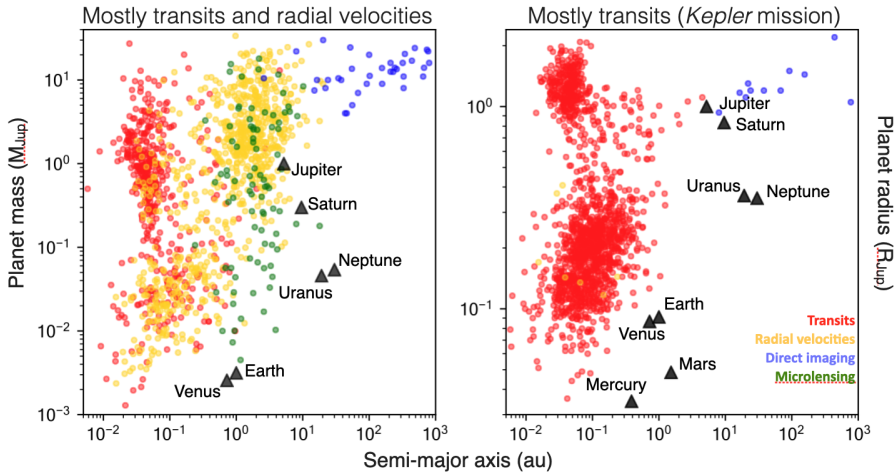


Figure 48: Confirmed exoplanet discoveries as of May 2020, with data obtained from the NASA Exoplanet Database (<https://exoplanetarchive.ipac.caltech.edu>). The left panel shows the distribution of masses and semi-major axes, whereas the right panel shows the planet radius versus semi-major axes. Exoplanets are represented with circles and colour-coded by their discovery method, while the planets of the Solar system are indicated with black triangles.

Interpreting these big Earths (or small Neptunes) becomes increasingly complicated, because in many cases it is not possible to determine from the size of the planet (more specifically from its mass and radius, giving us the bulk density) what these planets are made of. A degeneracy between interiors and atmospheres exist [1, 77], which makes it equally likely to have a big, light core with a thin atmosphere, or a smaller, denser core with a more massive extended atmosphere.

To complicate the picture even further, it is now clear that any types of planet is affected by the presence of aerosols [5, 6, 41, 46, 64, 86]. These could be clouds coming from condensation of species (some occurring even at the high temperatures of the hot Jupiters), or hazes created by photochemical processes, similar to those detected in the atmosphere of Titan, the biggest moon of Saturn.

Ultimately, if we want to assess the true nature of exoplanets, we need to access more information about them, and the best way we can do that observationally is by studying their atmospheres [57]. Broadly speaking, studying exoplanet atmospheres implies measuring their physical and chemical conditions to infer the planet structure (interior/envelope), the elemental abundances (related to formation and evolution scenarios), and the energy balance (how radiation is processed and heat transported). While we are still far from obtaining all this information in a coherent, robust, and complete fashion, observations have progressed enormously since the very first detection of sodium by [19]. This Chapter describes one particular observational technique based on spectroscopy at very high spectral resolution. To understand what this entails and how it is used, we will start by reviewing the fundamental concepts related to spectroscopy.

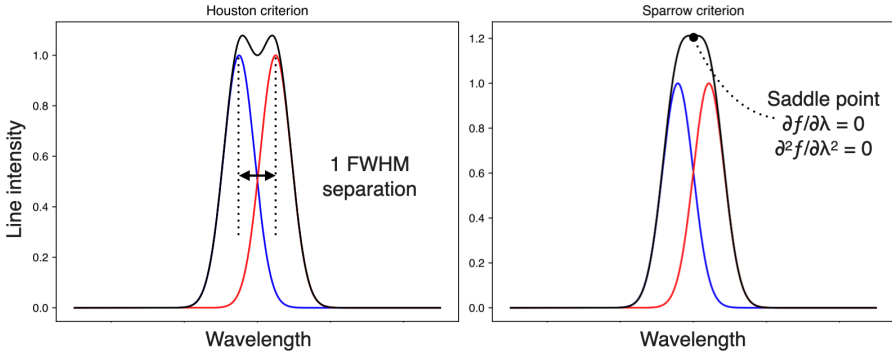


Figure 49: Illustration of the two criteria to define two spectrally-resolved lines, i.e. the Houston (left panel) and the Sparrow (right panel) criterion.

1.2 Resolution vs Resolving Power

In spectroscopy, *resolution* generically refers to the ability to distinguish photons of similar energy. Practically, this is often referred to as the ability to discriminate two narrow spectral lines with peaks separated by a difference in wavelength $\Delta\lambda$, and resolution is quantitatively defined as the minimum $\Delta\lambda$ at which the two lines are still distinguishable.

In this context we mention two main criteria that are adopted to define $\Delta\lambda$, that is the *Sparrow* and the *Houston* criteria. As shown in Figure 49, under the Sparrow criterion two lines are considered resolved when their combined profile has a saddle point, i.e. the first and second derivatives are both null. The Houston criterion, on the other hand, assumes that two spectral lines are resolved when their combined profile has two peaks separated by at least one full-width half-maximum (FWHM). Throughout this Chapter we will adopt the Houston criterion when referring to spectrally resolved lines.

A quantity closely related to resolution is *resolving power* R , that is

$$R = \frac{\Delta\lambda}{\lambda}, \quad (28)$$

i.e. the resolution divided by the wavelength of light. Resolving power has the more practical advantage that it can be considered nearly constant over the wavelength range of most spectrographs. Furthermore, R can be related to the radial velocity v of an astronomical object. In the classical limit ($v \ll c$) we have

$$R = \frac{\Delta\lambda}{\lambda} \simeq \frac{v}{c}, \quad (29)$$

where $c = 3 \times 10^8 \text{ m s}^{-1}$ is the speed of light. It follows that the minimum velocity shift Δv that can be resolved by a spectrograph is $\Delta v = c/R$.

In this Chapter we broadly separate the value of resolving power into three main regimes:

- Low-resolution spectroscopy (R up to 500): it results in $\Delta v \geq 600 \text{ km s}^{-1}$. This is generally too big to investigate the dynamics of most astrophysical sources, but still sufficient on the cosmic scales of galactic redshift.
- Medium-resolution spectroscopy ($R = 500\text{-}10,000$): it results in $\Delta v \geq 30 \text{ km s}^{-1}$. This allows us to start resolving broad spectral lines in astronomical spectra and derive physical properties from them.

Spectrograph	R	Telescope	Location
CRILES	100,000	VLT 8.2m	Cerro Paranal, Chile
CRILES+	100,000	VLT 8.2m	Cerro Paranal, Chile
CARMENES	80,000	CAHA 3.5m	Calar Alto, Spain
IRD	75,000	Subaru 8m	Mauna Kea, Hawaii
SPIRou	70,000	CFHT 3.5m	Mauna Kea, Hawaii
NIRPS	70,000	NTT 3.6m	La Silla, Chile
NIRSPEC	50,000	Keck 10m	Mauna Kea, Hawaii
GIANO	50,000	TNG 3.6m	La Palma, Spain

Table 3: List of near infrared high-resolution spectrographs currently available at ground-based telescope facilities, or planned for the near future. Indicated are the resolving power, the telescope name and its diameter, and the location of the observatory.

- High-resolution spectroscopy ($R = 10,000\text{--}200,000$): it results in Δv as small as 1.5 km s^{-1} . This allows us to explore the detailed shape of most spectral lines, and measure accurate motion of the objects generating these lines.

Needless to say, the boundaries between the various spectral regimes are quite arbitrary and possibly dependent on the discipline or application. In this Chapter we will focus on high resolution spectroscopy, and in particular on the infrared portion of the spectrum. There are several high-resolution spectrograph planned or currently in operation across the globe, and a non-exhaustive list can be found in Table 3.

1.3 Using spectroscopy to find exoplanets

If a star hosts a planet, it will orbit the common centre of mass of the system. Therefore, even if the planet is too faint and too close to the parent star to be spatially distinguished in the sky, we can still detect the indirect effect on the star itself. This is a periodic Doppler shift of the stellar spectral lines, with amplitudes up to a few hundreds m s^{-1} , but as low as 0.1 m s^{-1} for an Earth-Sun analogue. These radial velocities are well below the velocity resolution listed in Section 1.2 for current instrumentation, however a particular technique is used – called *cross correlation* – that allows us to increase the precision in radial-velocity measurements by combining the information from many lines in the stellar spectrum. We will extensively talk about cross correlation in Section 2.2.

The first exoplanet orbiting a main-sequence star was discovered around 51 Pegasi in 1995 [55] with the method outlined above, and this measurement granted the two authors of the discovery a Nobel prize in 2019. 51 Pegasi b is a Jupiter-size exoplanet orbiting its parent star in 3.5 days, and it causes a wobble in the spectral lines of 51 Peg of about 50 m s^{-1} . Whereas at the time of the discovery 10 m s^{-1} was considered an incredible technical achievement, most of the spectrographs in operation today are stable enough to deliver consistently a precision of $0.5\text{--}1 \text{ m s}^{-1}$. For example, in 2016 the closest star to the Sun (Proxima Centauri) was found to host an exoplanet with a minimum mass similar to the Earth [4]. The measured radial velocity amplitude of the parent star is only 1.4 m s^{-1} , which is equivalent to a slow walking pace!

Some spectrographs today are so stable that they can achieve an instrumental precision of 0.2 m s^{-1} over a night of observation. At this level, astrophysical signals start to dominate, first of all the radial velocity signal mimicked by stellar activity. Even the most accurate

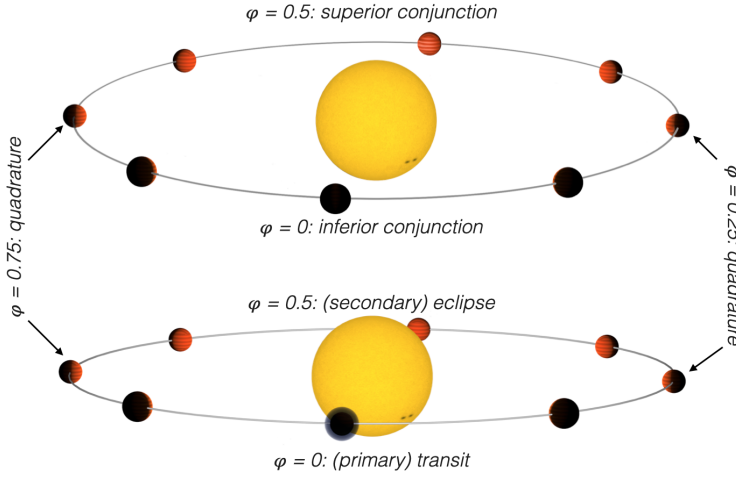


Figure 50: Definition of the orbital phase φ and orbital configurations for transiting (bottom drawing) and non-transiting (top drawing) exoplanet systems. Credits to E. de Mooij for the original drawing.

modelling to date struggles to push stellar RVs below the 0.5 m s^{-1} threshold, however to detect an Earth-Sun analogue we would need to hit the 0.1 m s^{-1} precision.

1.4 Exoplanets and their orbits

For the majority of this Chapter we will refer to exoplanets on circular orbits. The reason for this is that most of the observations focus on short-period planets (i.e. orbital periods at most of a few days), and these are thought to circularise on short time scales due to tidal interaction with their parent stars [70]. We will define *orbital inclination* i the angle between the line of sight and the normal to the plane of the orbit. So a transiting exoplanet (a so-called edge-on system) will have $i \approx 90^\circ$ while a face-on system will have $i \approx 0^\circ$.

In addition to the orbital inclination, a quantity that will be widely used is the *orbital phase* φ , that is the fraction of the orbit the planet has completed. We define the zero point in phase as the mid-transit (for a transiting planet), or more generally the point when the planet is closest to the observer, which is called inferior conjunction (see Figure 50). The centre of secondary eclipse (or superior conjunction) will thus have $\varphi = 0.5$ and the quadrature points $\varphi = 0.25$ or $\varphi = 0.75$.

If we assume a circular orbit of period P , it is straightforward to calculate the orbital phase as the fractional part of

$$\varphi = \frac{t - T_0}{P} \quad (30)$$

where T_0 is either the time of mid-transit or more generically the time of inferior conjunction (for a non-transiting system).

We can also compute the orbital velocity of an exoplanet as

$$v_{\text{orb}} = \frac{2\pi a}{P} \quad (31)$$

where a is the planet's semi-major axis, or simply the orbital radius for a circular orbit. Since we are seeing the system under an angle i , the radial velocity of an exoplanet while moving around the orbit can be found by projecting the orbital velocity along the line of sight:

$$RV_P = v_{\text{orb}} \sin(i) \sin(2\pi\varphi) = K_P \sin(2\pi\varphi), \quad (32)$$

where we have defined the semi-amplitude of the planet's radial velocity curve (i.e. the maximum projected orbital velocity) as K_P . As we will see in Section 3.1, this radial velocity is calculated in the reference frame of the exoplanet system. An observer on the Earth will measure additional time-varying radial velocity components.

1.5 Spectroscopy of transiting exoplanets

Observations of transiting exoplanets are made by carefully monitoring the light coming from the system as a function of time (commonly called a *light curve*) and looking for a periodic dimming of the starlight. The exact amount of dimming and its time evolution give insight on the period of the orbit, the relative size between star and planet, and the orbital inclination.

Measuring the stellar light accurately from the ground is challenging due to instrumental instability (flexure, pointing jitter, etc.) and the variable behaviour of the Earth's atmosphere (changing transparency due to water vapour, clouds, etc.). The common practice is to perform *differential photometry*, i.e. to measure the flux from the star of interest relative to one or more reference stars, possibly of similar brightness and spectral type. Assuming that systematic effects act similarly between target and reference, taking the ratio of light curves allows us to achieve a good correction of the unwanted effects. What is fairly straightforward in theory is only rarely achieved in practice, and this is why ground-based differential photometry still struggles to detect transits with depth much below 10^{-3} relative to the flux out of transit. While this is sufficient for giant to Neptune-size exoplanets orbiting solar-type stars, when it comes to detecting smaller transiting exoplanets space photometry is the only method of observation providing sufficient precision. This is why the most successful survey of transiting exoplanets – the *Kepler* mission [9] – was a spacecraft.

The relative geometry between a transiting exoplanet and the parent star can be exploited to access information about the planet's atmosphere, and it is schematically represented in Figure 51. In this section we will review the three main measurements that can be achieved with a particular focus on order-of-magnitude estimates of the planet-to-star flux ratios.

1.5.1 Transmission spectroscopy

When an exoplanet passes in front of the parent star – what is called a primary transit or more simply *transit*, a portion of the stellar surface is blocked by the opaque planet disk causing a corresponding dimming in the measured light from the system. If a planet possesses an atmosphere, it will selectively block light at certain frequencies corresponding to the energy transitions of the atoms and molecules in the gas phase. Increased absorption causes the planet to appear slightly bigger, therefore blocking slightly more starlight and producing a deeper transit (see inset in Figure 51 for a schematic representation of the phenomenon). By measuring the depth of a transit as a function of wavelength it is thus possible to build a *transmission spectrum* of an exoplanet, that is planet size versus wavelength.

In this Section we want to obtain a first-order estimate of the strength of atmospheric transmission signals. The depth D of an exoplanet transit is given by

$$D = \left(\frac{R_P}{R_\star} \right)^2, \quad (33)$$

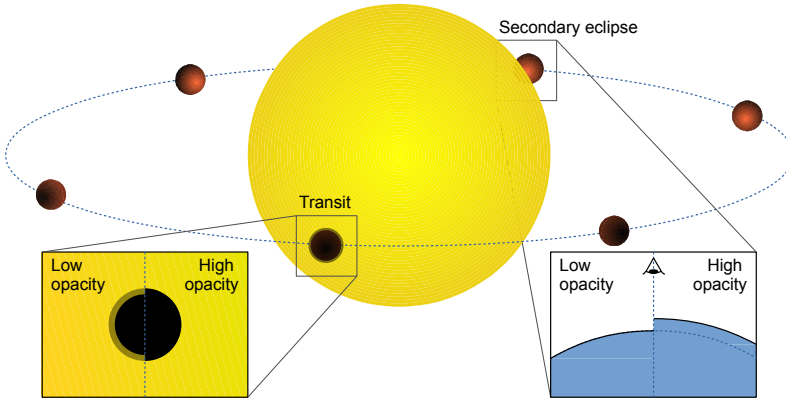


Figure 51: Relative geometry between a transiting exoplanet and its parent star, with highlights on the two principal configurations used to study exoplanet atmospheres, namely transmission and emission spectroscopy.

where R_P and $R_\star$ are the planet and stellar radii, respectively. For a Jupiter-size planet orbiting a solar-type star $D \sim 10^{-2}$, for an Earth-size planet around a solar-type star $D = 8 \times 10^{-5}$. In transmission spectroscopy we are interested in the change in transit depth with wavelength, that is $\Delta D(\lambda)$, which is approximately the ratio between the area of the atmospheric ring, $A_{\text{ring}}(\lambda)$ and the stellar surface area $A_\star$. If we assume that the planet's atmospheric ring has thickness δ , we obtain

$$\begin{aligned} A_{\text{ring}} &= \pi(R_P + \delta)^2 - \pi R_P^2 = \\ &= \pi(R_P^2 + 2R_P\delta + \delta^2 - R_P^2) = \\ &= \pi\delta(2R_P + \delta), \end{aligned} \quad (34)$$

where we have omitted the explicit dependence on wavelength λ . To estimate the order of magnitude of δ , we generally express it as a “number of scale heights”, that is $\delta = nH$. A scale height H is the change in altitude over which the planet's pressure changes by a factor of e (≈ 2.718) and can be expressed as function of physical parameters:

$$H = \frac{k_b T}{\mu g}, \quad (35)$$

where k_b is Boltzmann's gas constant, T is the temperature, μ the mean molecular weight and g the gravitational acceleration at the surface of the planet. The atmosphere of a typical gas giant, dominated in mass by molecular hydrogen ($\mu = 2.23 \text{ g mol}^{-1}$), will have approximately 8 times bigger signals than a hypothetical planet dominated by water ($\mu = 18.0$).

How many scale heights should we consider? The value of n is connected to the change in opacity κ as a function of wavelength. More precisely, $n \approx \log \Delta\kappa$. Since κ can change up to 2 orders of magnitude between the centre of a strong molecular line and the regions far from molecular absorption, we can expect n up to $\log(100) \approx 5$. In the centre of strong atomic lines, such as the sodium doublet in the optical, n can be as high as 10.

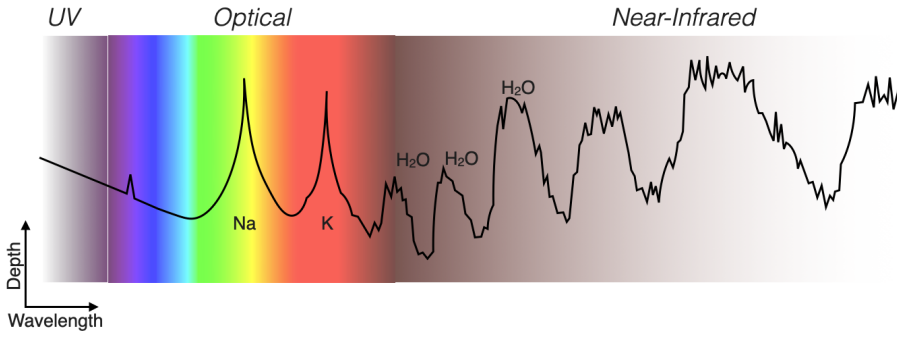


Figure 52: Schematic representation of the transmission spectrum of a cloud-free hot Jupiter, showing the strong alkali lines in the optical and molecular bands from water vapour in the near infrared.

Using our definition of δ into Equation 34 we obtain:

$$A_{\text{ring}} = \pi n_{\lambda} H (2R_p + n_{\lambda} H) \simeq 2\pi n_{\lambda} R_p H, \quad (36)$$

where we have neglected terms of $o(H^2)$ and explicitly indicated the wavelength dependence of n . We can therefore estimate the change in transit depth as

$$\Delta D = \frac{2\pi n_{\lambda} R_p H}{\pi R_{\star}^2} = \frac{2n_{\lambda} R_p H}{R_{\star}^2}. \quad (37)$$

For a hot, giant exoplanet ($T = 1,500$ K, $\mu = 2.23$), $\Delta D \sim 100$ ppm, while for a cooler “Super-Earth” ($R_p = 2.5R_{\oplus}$, $T = 800$ K) with an atmosphere dominated by water vapour ($\mu = 18$) this is only $\Delta D \sim 4$ ppm. Here we have introduced the concept of *part per million* (ppm) which is a ratio of 10^{-6} and it is useful to express the small signals expected from exoplanet atmospheres.

Interpreting transmission spectra can reveal a wealth of information about the planet’s atmosphere. Figure 52 shows a schematic representation of a typical transmission spectrum of a hot-Jupiter in the absence of clouds and for a solar composition (metallicity and C/O). At optical wavelengths, the spectrum is dominated by the strong features of alkali lines, the Na and K doublets, and by Rayleigh scattering from molecular hydrogen giving rise to a slope towards the UV. In the red part of the optical, some weak molecular lines from water vapour start to appear. For very hot giants, there can be additional molecular absorption from TiO or VO in the gas phase (not shown in the Figure). Molecular bands become more prominent at infrared wavelengths, with the spectrum dominated by the main molecular bands of water vapour. Other species such as methane, carbon monoxide, carbon dioxide, hydrogen cyanide, acetylene, ethylene, and ammonia also contribute to the chemical network determining the abundance of species in equilibrium. However, due to the dominant opacity from water vapour, these species are generally thought to be less readily detectable, especially at low resolution. They can play an important role if the elemental abundances of elements (in particular oxygen versus carbon) is significantly higher than the solar value of C/O = 0.55 [51] or if disequilibrium chemistry cannot be neglected [59].

In [86], Extended Data Figure 4 shows the effect of hazes, clouds, and metallicity on the shape of the transmission spectrum. *Hazes* are aggregates mostly produced by photochemical processes (as opposed to clouds that are the product of condensation) and they

increase the opacity of the planet atmosphere at all wavelengths, muting the spectral features from atoms and molecules. In addition, they have a very strong Rayleigh scattering signature in the optical spectrum, greatly enhancing the slope of the spectrum towards the blue. *Clouds* also increase the opacity at all wavelengths, however they behave more uniformly with wavelength, which means they give rise to flat, heavily “muted” spectra. Finally, increasing metallicity has the main effect to increase the mean molecular weight of the atmosphere, thus reducing the scale height and therefore the overall amplitude of the whole spectrum.

It is worth noting that transmission spectra are plotted as planet radius (or transit depth, or number of scale heights) versus wavelength, so increased absorption appears as a positive deviation of the spectrum, which can be somewhat confusing when thinking to absorption producing dips in a normal spectrum (flux versus wavelength).

[86] also present a comparative study of 10 exoplanets with high-quality space observations in the optical and NIR obtained with HST STIS and WFC3. Although their interpretation is still open to debate, they advocate that the different strength of the water bands in the NIR it is not due to changing abundance of the species, but rather to the muting effects of clouds or hazes. However, other comparative studies [5, 64] claim that even when considering clouds there is still evidence for lower water abundance than expected from equilibrium chemistry and abundances consistent with the solar value. This discrepancy highlights how difficult it is to discriminate these scenarios from low-resolution observations alone, given the extreme precision (a few tens of ppm) required from the observations.

1.5.2 Information from secondary eclipses: reflected light

When a planet passes behind the parent star, we measure an *eclipse* (sometimes referred to as secondary eclipse in exoplanet science). This is a direct measurement of the missing planet flux, which at optical wavelengths is typically dominated by reflected starlight. Note that this is not true for every planet, as particularly hot giants can still show a significant thermal emission in the red optical, and therefore it is somewhat difficult to interpret the eclipses of these systems.

A first-order estimate of reflected light signals can be obtained by writing the stellar flux received by a planet orbiting at distance a from a star with luminosity $L_\star$:

$$F_{\star,P} = \frac{L_\star}{4\pi a^2}. \quad (38)$$

The above represents the stellar luminosity redistributed over a sphere of surface $4\pi a^2$. The planet reflects back a fraction of the incident stellar flux depending on its reflectivity, which is called *albedo* (A). The amount of energy reflected back will therefore be

$$E_{\text{refl}} = A\pi R_p^2 F_{\star,P} = \frac{AR_p^2 L_\star}{4a^2} \quad (39)$$

In writing the above equation we have assumed the planet is a disk of area πR_p^2 . An observer at distance d from the system will therefore measure a flux ratio between the planet and the star equal to

$$\frac{F_{P,\text{refl}}}{F_\star} = \frac{AR_p^2 L_\star}{(4\pi d^2)(4a^2)} \frac{4\pi d^2}{L_\star} = A \left(\frac{R_p}{2a} \right)^2. \quad (40)$$

Equation 40 suggests that the best reflected-light signals should be measured for giant planets very close to their parent stars, however theoretical [94] and observational [26] evidence show

that hot Jupiters are poorly reflective at optical wavelengths. This is due to the presence of Na and K atoms, which absorb incoming radiation very effectively across a broad wavelength range, due to their strong wings extending away from the centres of absorption lines. As an example, the exoplanet WASP-33 b ($a = 0.025$ au, $R_p = 1.6R_{\text{Jup}}$) has an optical eclipse depth of only 20 ppm, corresponding to an albedo of 0.1. If we assume this value as typical, an average hot Jupiter with $a = 0.04$ au and $R_p = 1.15R_{\text{Jup}}$ would only show a reflected signal of 5 ppm. Due to the quadratic dependence on planet radius and semi-major axis, signals quickly scale down for terrestrial temperate planets. The Earth around the Sun ($A = 0.35$) would have a reflected light signal of only 0.16 ppb (1 ppb = 10^{-9}).

It is therefore not surprising that measurements of reflected starlight are incredibly rare, and so far limited to a single broad-band photometric filter from space missions such as Kepler [9] and TESS [73]. There is one notable exception, that is the wavelength-resolved albedo spectrum of HD 189733 b [27].

1.5.3 Information from secondary eclipses: thermal emission

At infrared wavelengths, the secondary eclipse depth is a direct measurement of the missing planet thermal emission, and therefore measuring the eclipse depth as a function of wavelength allows us to build the emission spectrum of an exoplanet. This is also called *day-side spectroscopy* in the context of exoplanets. The principle behind the method is that at wavelengths where the exoplanet atmosphere is more opaque, the observer sees radiation emitted from layers higher up in the planet's atmosphere, as shown by the inset in Figure 51. According to whether the planet has a temperature decreasing or increasing with altitude, the corresponding layer will appear darker (absorption) or brighter (emission) compared to the deeper layers, the latter generally referred to as the *continuum*. As we will see a bit more in detail in Section 2.5, this principle is at the root of whether we see emission or absorption lines in the planet's day-side spectrum, and it is exploited by HRS to recognise these two scenarios.

Due to the fact that thermal emission is dependent on opacity, and that the continua of planet and stars are dependent on both temperature and wavelength, we can only estimate the emission signals of an exoplanet by explicitly considering the wavelength dependence. We will approximate star and planet as black bodies at temperatures T_{eq} and T_{eff} respectively, and use Planck's formula for black-body radiation, that is

$$B(\lambda, T) = \frac{2hc^2}{\lambda^5} \frac{1}{\exp[hc/(\lambda kT)] - 1}, \quad (41)$$

where h is Planck's constant, c the speed of light, k Boltzmann's constant. Note that B has units of [Energy / time / area / wavelength / solid angle], but in the approximation of isotropic radiation the integral over the solid angle trivially amounts to a factor of π , so flux can be obtained via $F = \pi B$ and luminosity via $L = 4\pi R^2 B$.

We will now estimate the planet's temperature T_{eq} . We will neglect any internal source of energy, which is a good assumption for evolved exoplanets that have long completed their cooling cycle since formation. In this scenario, the only energy available is the incoming stellar radiation. We can use similar arguments as for the treatment of reflected light in Section 1.5.2 to compute the total energy absorbed per unit time:

$$E_{\text{abs}} = (1 - A)\pi R_p^2 \frac{4\pi\sigma R_\star^2 T_{\text{eff}}^4}{4\pi a^2}, \quad (42)$$

where we have accounted for the planet albedo and written the stellar luminosity via the Stefan-Boltzmann formula for a black body, with σ being the Stefan-Boltzmann constant.

The energy re-emitted by the planet as a black body will be

$$E_{\text{out}} = 4\pi\sigma R_p^2 T_{\text{eq}}^4. \quad (43)$$

Equalling the incoming and outgoing energies leads to an expression for the planet's temperature called the equilibrium temperature:

$$T_{\text{eq}} \simeq T_{\text{eff}} \sqrt{\frac{R_\star}{a}}, \quad (44)$$

where we have neglected any additional scaling factor. The exact value of this factor depends on the properties of the exoplanet atmosphere, in particular on its efficiency to redistribute the incoming stellar radiation both in latitude and longitude. For an analytical treatment of some of the main cases please see [23].

We can now put all the information together to estimate the flux ratio between the thermal emission of an exoplanet and its parent star:

$$\frac{F_{\text{P,em}}(\lambda)}{F_\star(\lambda)} = \frac{L_p(\lambda)}{4\pi d^2} \frac{4\pi d^2}{L_\star(\lambda)} = \frac{B(T_{\text{eq}}, \lambda)}{B(T_{\text{eff}}, \lambda)} \left(\frac{R_p}{R_\star}\right)^2. \quad (45)$$

A hot, giant exoplanet (1,500 K) would have a flux ratio of 30 ppm at 1.2 micron, rising to 290 ppm at 2.3 micron, a more comfortable value than reflected light.

Due to the strong dependence on temperature, however, a warm Neptune with 800K would have the above values reduced to 3 ppb and 0.64 ppm, respectively, for the two wavelengths considered.

1.5.4 Information from the whole orbit: phase curves

The exquisite precision of space photometry allows us to study not only the decrease in starlight during transit and secondary eclipse, but also the periodic variations while the planet moves around the orbit. These light curves are called *phase curves* and they allow us to estimate the changing contribution from the day- and night-side of the planet as seen by the observer. Once again, the interpretation of these phase curves is different whether they are obtained at infrared wavelengths (dominated by thermal emission and thus mapping the changing temperature between the planet's day- and night-side) or in the optical (dominated by reflected light and/or indirect effects such as tidal deformation of the parent star or Doppler beaming of stellar radiation). In terms of constraining the properties of the exoplanet atmosphere, the most stringent measurements are infrared phase curves, as they are directly related to the efficiency of energy redistribution from the sub-stellar point to the night-side. Early measurements of phase curve, for instance, allowed astronomers to verify one fundamental model prediction, that is hot Jupiters develop strong equatorial winds towards the east (similar to jet-stream currents on Earth) capable of transferring significant amount of energy [84]. Experimentally, this is verified by the fact that thermal phase curves do not peak at phase 0.5 (i.e. when the planet's sub-stellar point faces the observer), but slightly earlier [42].

Even from space, switching from photometric phase curve (i.e. integrated over the pass-band of a filter) to spectroscopic phase curves (resolved as a function of wavelength) is a major technological endeavour, so far limited to a handful of objects (at least pre-JWST). The most detailed study to date using spectroscopic phase curves is on the atmosphere of WASP-43 b [93].

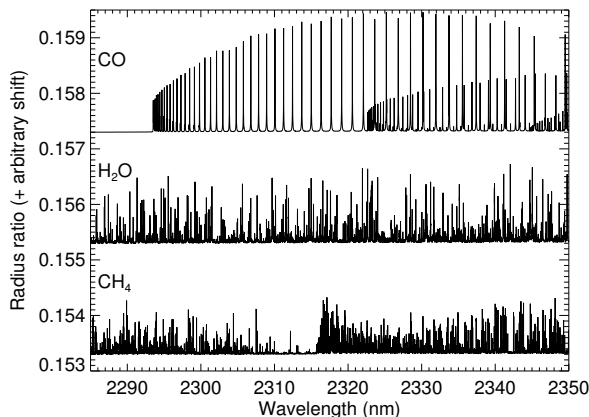


Figure 53: High resolution transmission spectrum of the hot Jupiter HD 189733 b computed around $2.3 \mu\text{m}$ for three important molecular species, namely carbon monoxide (top), water vapour (middle), and methane (bottom). The three spectra are shifted vertically for clarity.

2 High-resolution spectroscopy of exoplanets

2.1 Key characteristics

High resolution spectroscopy at visible and near-infrared wavelengths ($0.5\text{--}3.5 \mu\text{m}$) is only possible from the ground at the moment. This means that high-resolution spectra will be imprinted not only by the spectrum of the astrophysical target, but also by the spurious transmission spectrum of the Earth's atmosphere. We will refer to this spectrum as the *telluric spectrum*.

If we compute model transmission spectra of an exoplanet at increasingly higher resolution, molecular bands start being resolved into their individual lines. Energy levels in molecules depend on the specific structure and degrees of freedom, thus each species has a unique sequence of lines, a sort of fingerprint, that is well distinguishable at very high resolving powers. This is shown in Figure 53 for three main species around 2.3 micron, namely CO, H₂O, and CH₄. Species can be thus identified by matching models to data, commonly done through the technique of cross correlation as described in Section 2.2.

A second peculiar consequence of observations at high resolving power is that the orbital motion of exoplanets is also detectable. For a typical hot Jupiter on a circular orbit, orbital velocities exceed 100 km s^{-1} . This means that over a few contiguous hours of observations the planet's radial velocity changes by tens of km s^{-1} , especially when the planet is observed near conjunction. In contrast, the telluric spectrum and in first approximation the stellar spectrum are nearly stationary in wavelength. This allows us to robustly disentangle the planet from any contaminants in velocity space, and also to directly measure the planet's radial velocity. We will see in Section 2.3 how this can be used to constrain the planet mass and orbital inclination. In Section 3.1 we will relax the approximation that the stellar spectrum is stationary in Doppler velocity and discuss some important consequences for the correct interpretation of measurements.

The changing Doppler signature of exoplanet spectra is schematically represented in Figure 54. This is a schematic model showing a small portion of the spectrum as a function of planet orbital phase (recall our convention about orbital phases explained in Section 1.4).

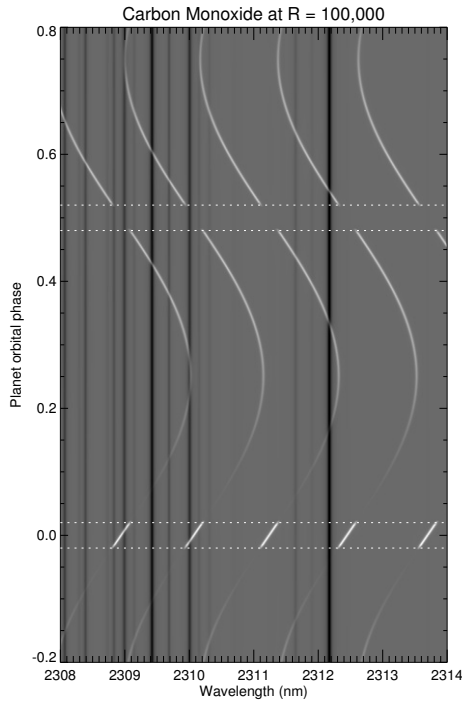


Figure 54: Diagram showing the change in Doppler shift experienced by the spectral lines of a transiting exoplanet (in white) versus the stationary absorption caused by the Earth’s atmosphere (in black). It shows that the signal is maximised during transit (around $\varphi = 0$) or just before or after secondary eclipse ($\varphi \sim 0.5$).

Represented in white is carbon monoxide absorption in the planet spectrum, which is subject to a stark change in Doppler shift. In contrast, telluric lines (in black) are completely stationary. It is clear that the best exoplanet signals are obtained during transit and around secondary eclipse. Differently from emission spectroscopy explained in Section 1.5.3, we are not comparing the fluxes in and out of eclipse, but rather we are targeting the planet’s spectrum directly. This is an important distinction to make: with photometry or low-resolution spectro-photometry time resolution is used to compare different parts of a light curve (in- or out- of transit/eclipse), while at high spectral resolution time resolution is used to allow the planet’s radial velocity to change. Planet and stellar/telluric spectra are thus discriminated in Doppler (or velocity) space, rather than in flux. An important corollary is that we can observe the high resolution emission spectrum of non-transiting planets, as we will further see in Section 2.3. As of today, this is the only method capable of investigating the atmospheres of evolved, non-transiting planets, and it has the benefit that it allows us to also measure the planet’s mass and orbital inclination, thus breaking a fundamental degeneracy of non-transiting exoplanet systems.

2.2 The signal-to-noise formula and cross correlation

With the exception of direct imaging techniques, which are only marginally relevant for future applications of high-resolution spectroscopy (see Section 5.2), exoplanet observations do not

distinguish between photons from the star and those from the planet, i.e. star and planet are not spatially separated. We therefore measure photons coming from both sources, but stars are orders of magnitude brighter than planets, which has a profound impact on the main sources of noise of our measurements.

Before trying to estimate the signal-to-noise ratio (S/N) of these measurements, let us remind ourselves that photons obey Poisson statistics, that is the standard deviation of a measurement is $\sigma = \sqrt{N}$, where N is the total number of photons counted. The *signal* in our case is the total number of photons emitted, absorbed, or reflected by the planet, which we indicate with N_p . The *noise* budget includes all the measured photons instead, that is N_p plus the photons from the parent star $N_\star$. Since $N_\star \gg N_p$, the simplest S/N formula that we can write is therefore:

$$\frac{S}{N} = \frac{N_p}{\sqrt{N_\star + N_p}} \approx \frac{N_p}{\sqrt{N_\star}}. \quad (46)$$

Going into more detail, there are other sources of noise increasing the denominator. Firstly, a measurement contains additional background photons, due e.g. to scattering or emission in our own atmosphere. We will call this term N_{sky} . Secondly, detectors are not perfect photon counters and they produce two main additional sources of error, the former being additional dark current (N_{dark} electrons generated because of the non-zero temperature of the electronics), and the latter being a constant read-out noise per pixel, σ_{RO} , due to the electronics “counting” the total number of electrons recorded on the detector.

Due to the fact that all these sources of noise are independent, the S/N formula can be trivially generalised by summing them in quadrature inside the square root at the denominator:

$$\frac{S}{N} = \frac{N_p}{\sqrt{N_\star + N_{\text{sky}} + N_{\text{dark}} + n_{\text{pix}} \sigma_{\text{RO}}^2}}, \quad (47)$$

where n_{pix} is the number of pixels co-added to extract a spectrum.

Let us consider an observation of a bright star and therefore use Equation 46, neglecting the additional sources of noise. We can manipulate this equation by multiplying and dividing by $N_\star$ and write an explicit dependence on the planet-to-star flux ratio:

$$\frac{S}{N}(\lambda) = \frac{N_{\lambda,p}}{N_{\lambda,\star}} \sqrt{N_{\lambda,\star}}. \quad (48)$$

To allow for a wavelength-dependent S/N, we have also explicitly indicated the quantities dependent on λ . Even for the strongest transmission spectroscopy signals, the planet/star flux ratio ($N_{\lambda,p}/N_{\lambda,\star}$) is on the order of 10^{-3} , while at very high spectral resolution on 8-meter class telescopes even the brightest stars hardly deliver more than $N_{\lambda,\star} \sim 10^6$ photons recorded over a few hours of observations on each individual spectral channel. This means that a single feature of the planet’s spectrum, e.g. one spectral line, would be detectable at most at a S/N = 1. It follows that some technique is needed to “amplify” the planet’s signal. One of such techniques is called *cross correlation*.

As the name suggests, the cross correlation is a mathematical operator that measures the level of correlation between two vectors, in our case a noisy spectrum and a model spectrum, as a function of a shift, which is technically called *lag*. Figure 55 shows an example of a CO model spectrum with noise added such as each spectral line has S/N of 5 (top panel) and 0.8 (bottom panel). The latter value is representative of a typical observation of CO with CRIRES at the Very Large Telescope, e.g. [16], and while the spectrum is completely buried in the noise in this case, its cross correlation with the noiseless model clearly peaks at zero lag. To understand what is happening to the S/N formula (Equation 46) it is important to realise that

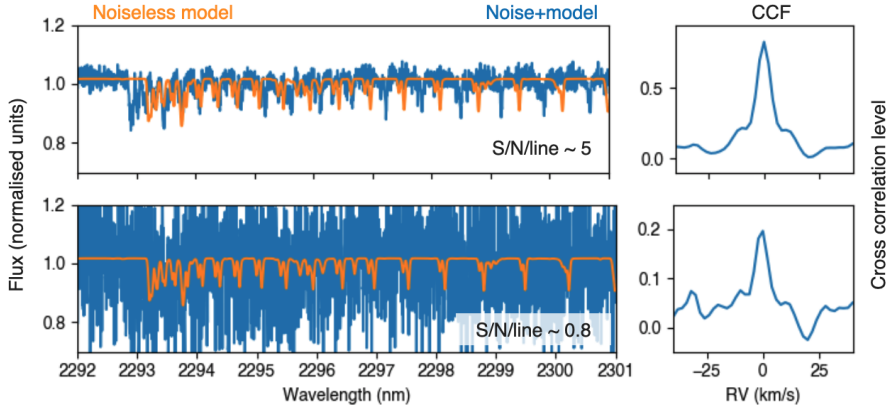


Figure 55: Illustration of the principles of cross correlation. The left column shows the comparison between a noiseless model of CO (in orange) and the same model with noise added (in blue) at two values of signal-to-noise ratio (S/N) per line, namely 5 (top) and 0.8 (bottom). The right column shows the resulting cross-correlation function (CCF) as a function of the relative shift between the noiseless and noisy spectra. It shows that while at S/N = 0.8 it is impossible to distinguish any spectral lines, the CCF correctly peaks at zero lag and significantly above the noise floor, due to the co-added signal of many CO lines.

the cross-correlation acts as a sort of matching filter, thus co-adding the signal from multiple spectral lines. If n lines are co-added, N_p and $N_\star$ both increase by a factor of n :

$$\left(\frac{S}{N}\right)_{\text{CCF}} = \frac{nN_p}{\sqrt{nN_\star}} = \sqrt{n} \left(\frac{S}{N}\right)_{\text{line}}. \quad (49)$$

However, due to the fact that photons obey Poisson statistics, while the signal increases by a factor of n , the noise only increases by a factor of $\sqrt{n}$, thus delivering a net gain of $\sqrt{n}$ in S/N. Referring again to the VLT/CRIRES observations of a typical hot Jupiter, the CO band around $2.3 \mu\text{m}$ has about 50 strong CO lines, meaning that the total cross correlation will have a S/N of $(0.8 \times \sqrt{50}) \sim 6$, resulting in a strong detection. More complex molecules, such as water vapour and methane, can have thousands of strong lines across the near infrared, providing a much higher amplification factor.

2.3 Understanding the analysis of high-resolution exoplanet spectra

The idea behind the use of high resolution spectroscopy to probe exoplanet atmosphere is actually older than the first space observations [20, 22]. However, it took another decade to develop suitable instrumentation (stable and with a wide-enough spectral range) and most importantly a suitable data analysis technique. The latter is truly peculiar to the method and therefore deserves a whole section to cover its principles. Here we will review the analysis of near-infrared data, which is somewhat more complex than optical HRS. This is due to the often negligible telluric absorption in the visible, allowing the extension of techniques suitable for photometry (e.g. taking the ratio between spectra in-transit and those out-of-transit) to high-resolution spectra [99]. The other important difference is that optical high-resolution spectroscopy typically analyses individual atomic lines (e.g. sodium and potassium) and therefore no cross correlation is used to amplify the planet's signal.

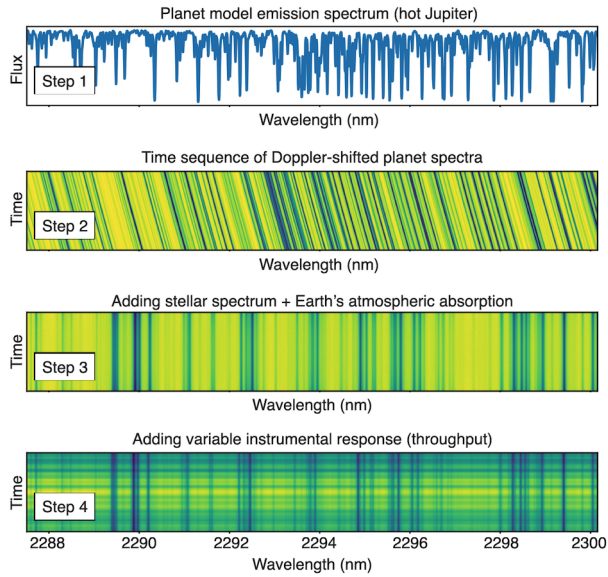


Figure 56: Illustration of the various spectral components present in a time sequence of HRS spectra. The emission spectrum of an exoplanet (Step 1) is subject to a Doppler shift variable by tens of km s^{-1} in a few hours of observations (Step 2). This spectrum is overwhelmed by the strength of the stellar spectrum mixed with it and the absorption lines in the Earth's atmosphere (Step 3), and additional instability causes the overall recorded flux to vary as a function of time (Step 4).

To start with, we will focus on the way data is organised and how the spectral components from the planet, the star, and the Earth's atmosphere look like. We will take an emission spectrum as example, but similar reasoning is applicable to transmission spectra. Step 1 in Figure 56 shows a model emission spectrum of an exoplanet containing water and carbon monoxide lines. Relative to the continuum of the star, the spectral lines in such models have depth of 100 ppm. During an observing night, a sequence of such spectra is observed at the telescope (typically a few tens of spectra each exposed for a few minutes at most), and spectra are re-organised in two-dimensional matrices with wavelength on the horizontal axis and time (or orbital phase) on the vertical axis. Due to the changing orbital radial velocity of the planet (non negligible over a few hours of observations), the planet's spectrum is subject to a changing Doppler shift, which is why the sequence of planetary spectra appears shifting, giving rise to slanted lines (Step 2). As explained in Section 2.2, in a real observation we do not separate the planet and stellar photons, so the faint planet spectrum gets completely overshadowed when stellar and telluric spectrum are added (Step 3). The depth of stellar and telluric lines is several tens of percent (and can get as high as 100% for very strong lines). Furthermore, due to the fact that the telluric lines can be considered stationary in the reference frame of an Earth's observer and stars also move at most by a few hundreds of m s^{-1} (see Section 3.1), the position of these lines does not change over the observing night, i.e. they fall on the same pixels of the matrix (along the spectral axis, or the x -axis in the Figure). Even though the telluric spectrum is stationary in wavelength, the level of absorption varies with time because of changes in the overall transparency (due e.g. to variable water vapour, clouds, etc.), and changes in the altitude of the target over the horizon (sometimes referred

to as geometric airmass). Furthermore, the pointing of telescopes is subject to small errors, meaning that the target star is tracked imperfectly leading to variation in the light entering the instrument. As a result, each observed spectrum has a slightly different number of photons recorded (typical variations on a good night are 20-30%), and this effect appears as a series of horizontal stripes (Step 4).

The analysis of high resolution infrared spectra is designed to “reverse-engineer” this process, that is to start with Step 4 in Figure 56 and recover the “cleaned” sequence of exoplanet spectra (Step 2). Since its first successful application by [91], several techniques have been developed to achieve this result, broadly divided into two categories: blind algorithms such as PCA [24, 66] or Sysrem [7], and de-trending by physical parameters [16, 80]. While operating similarly on the matrix of observed spectra, these two categories are conceptually very different in that the former focuses on a mathematical decomposition of a matrix in a series of eigenvectors, while the latter tries to estimate these eigenvectors from measurable quantities and physical parameters. In this context we will explain the latter category as it is the most immediate to understand:

1. Each spectrum (each row in Figure) is divided by its median count level. This removes variations in pointing and broad-band variations due to airmass. The resulting matrix is thus normalised (e.g. the continuum is approximately equal to 1).
2. The depth of telluric lines changes with airmass, so the next step involves fitting each telluric line (each column in the data) as a function of geometric airmass, and dividing through the fit.
3. Water vapour absorption departs from pure airmass dependence because it also depends on variable humidity. This produce visible residuals in panel. It is thus necessary to apply a second-order correction to water telluric lines, typically by fitting each data column by a low-order polynomial.
4. Generally the previous steps do an excellent job at normalising a large fraction of the data, however there are still some residual noisy columns due to imperfect telluric correction, saturated spectral lines, or damaged pixels on the detector. These few columns (at most a few percent of the total) are masked and not used further in the analysis.

As estimated in Section 2.2, even after the spectra are cleaned from the stellar and telluric absorption, the planet’s signal is too weak to be detected at this stage. We therefore apply cross correlation, i.e. we measure the level of correlation between the data and a set of models as a function of radial velocity. Since each spectrum of the sequence is correlated against the same radial velocity lag, we produce a matrix of cross correlation functions (CCFs), that is the level of correlation between data and models as function of radial velocity (on the horizontal axis) and time (on the vertical axis), see Figure 57. The planet spectrum will appear as an excess correlation (darker colours in the figure) at certain values of radial velocity, shifting in time due to the planet’s orbital motion (recall our explanation in Section 2.1). In real applications even at this stage it is hard to visually distinguish the excess correlation coming from the planet, however it is possible to further enhance the signal to noise by co-adding the cross correlation functions in time, i.e. along the sequence of planet radial velocity. We have seen in Equation 32 how in the stellar reference frame the planet’s radial velocity depends on the orbital phase φ and the semi-amplitude K_p . We will now add the motion of the system hosting the exoplanet relative to the solar system, which is called the *systemic velocity* v_{sys} , to obtain the general equation:

$$RV_P = v_{\text{sys}} + K_P \sin(2\pi\varphi). \quad (50)$$

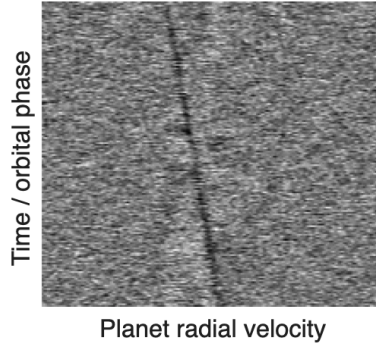


Figure 57: Simulated cross correlation matrix as a function of radial velocity (horizontal axis) and time (vertical axis) obtained by adding a strong ($20\times$ nominal) model of CO to real observations of exoplanet τ Boötis b ([16]). It clearly shows the excess correlation (darker values in the image) tracing the radial velocity of the planet, which is changing due to the varying orbital motion of the planet.

Generally for transiting exoplanets we know quite accurately the orbital velocity and inclination (determining K_P), the orbital phase, and the systemic velocity is also known for bright stars to a fraction of a km s^{-1} . For non-transiting systems, depending on how recent the last observations were and how strong is the evidence for zero eccentricity, φ might be known to a percent level, but the orbital inclination is unknown. As a consequence, the co-adding of the cross correlation functions is done as a function of two velocities, namely v_{sys} (determining a global shift of the CCFs), and K_P , changing the amplitude – and thus the slope – of the slanted line in Figure 57.

Figure 58 shows the total cross correlation signal as a function of the two velocities for the first non-transiting system successfully observed via HRS, namely τ Boötis b [16]. The detection in this particular case is obtained with CRIRES ($R = 100,000$) at the ESO Very Large Telescope by co-adding about 50 strong CO lines around $2.3 \mu\text{m}$. The planet was observed for a total of 15 hours on 3 different nights, covering orbital phases pre-, during, and post-superior conjunction, and resulting in a good range of radial velocities for the planet.

Once the planet's RV semi-amplitude K_P is determined, it is possible to calculate the planet's mass with methods akin to spectroscopic stellar binaries. If we denote with $K_\star$ the semi-amplitude of the stellar radial-velocity curve, the mass ratio of the planet is

$$\frac{M_P}{M_\star} = \frac{K_\star}{K_P}, \quad (51)$$

and therefore the planet mass is $M_P = M_\star K_\star / K_P$. It is also possible to estimate the orbital inclination of the system via

$$i = \sin^{-1} \left(\frac{K_P}{v_{\text{orb}}} \right), \quad (52)$$

where v_{orb} is the orbital velocity as defined via Equation 31. Table 4 lists the non-transiting planets for which masses and inclinations have been measured through the application of HRS.

Even for transiting planets, despite the good prior knowledge of φ , K_P , and v_{sys} , there is an advantage in performing the co-adding of the CCFs as a function of the two velocities. It ensures that the planet's signal is correctly detected at the expected radial velocity, and that

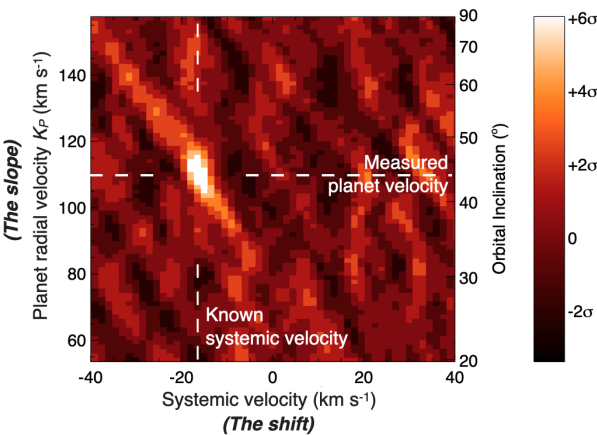


Figure 58: Detection of carbon monoxide absorption in the day-side spectrum of exoplanet τ Boötis b, observed at 2.3 micron with VLT/CRIRES ([16]). The dashed white lines show the known systemic velocity (vertical line) and the derived planet radial-velocity amplitude (horizontal line). The annotations on the two axes connect the quantities represented with the slope and the shift of the slanted line of excess correlation shown in Figure 57.

Planet	Inclination (degrees)	Mass (M_{Jup})	Reference
τ Boötis b	45.5 ± 1.5	5.95 ± 0.28	[16]
51 Pegasi b	> 79.8	0.46 ± 0.02	[14]
HD 179949 b	$66.2^{+3.7}_{-3.1}$	$0.96^{+0.04}_{-0.03}$	[98]
HD 88133 b	15^{+6}_{-5}	$1.02^{+0.61}_{-0.28}$	[66]
ν Andromedae b	24 ± 4	$1.70^{+0.33}_{-0.24}$	[67]
HD 102195 b	> 72.5	0.46 ± 0.03	[34]

Table 4: List of non-transiting exoplanets for which mass and orbital inclination were determined through near infrared high-resolution spectroscopy. Orbital inclinations compatible with transits were given as a lower limit. All the quoted error-bars are $1\text{-}\sigma$.

there is no other spurious detection in a different region of the parameter space. As we will see further on in this Chapter, there are spurious sources of correlated noise that survive the data analysis and might cause spurious detections elsewhere in the parameter space, so this sanity check is far from redundant.

2.4 Molecular and atomic species detected

As of May 2020, HRS has detected five molecular species: CO [11, 14, 16, 17, 24, 35, 74, 75, 81, 90, 91], H₂O [3, 7, 8, 10–12, 14, 17, 34, 35, 47, 66–68, 79, 81], HCN [17, 35], CH₄ [34], and TiO [61]. These species were all detected in the near-infrared with the exception of TiO, and they all share the conceptual processing of the data explained in Sections 2.2 and 2.3.

In addition, the higher stability of optical data and the relatively low influence of telluric lines allowed the extension of techniques originally designed for low-resolution spectro-

photometry to high-resolution transmission spectroscopy to extract the excess absorption in the core of strong atomic lines. This allowed us to detect atomic species such as H [38], He [2, 60], K [39], Ca [96], Na [72, 88], Ti and Fe [37].

In terms of exoplanet sample, a total of 25 objects were investigated at high resolution, among which 8 non-transiting planets, 15 transiting planets, and 2 directly-imaged planets. Contrarily to low-resolution studies, mostly reliant on information coming from the analysis and interpretation of light curves obtained with HST/STIS, HST/WFC3, and the 4 Spitzer photometric filters, the data and processing of HRS is highly heterogeneous, preventing so far comparative studies where the information from the whole sample is put together in a coherent fashion. Such comparative studies are instead available at low spectral resolution [5, 64, 86].

2.5 Thermal inversion layers

One key scientific question that can be answered trivially at high spectral resolution is whether exoplanet atmospheres show strong inversion layers. Originally proposed for planets with temperatures above 2,000 K (i.e. above the condensation limits of TiO and VO), thermal inversion layers have been controversial for many years and the source of much speculation [29, 43, 52, 62, 92]. Recently, evidence from low-resolution spectroscopy has suggested that there is indeed a class of hot giant exoplanets exhibiting thermal inversions, but there have significantly higher temperatures than initially thought ($T > 2,500$ K).

But what is an inversion layer in the first place? As spectral lines correspond to specific energy transitions in atoms and molecules, photons are absorbed much more effectively in the core of spectral lines than away from them. This means that line cores are formed higher up in the atmosphere, because an increased opacity will raise the level of the planet photosphere (recall the inset in Figure 51). In emission spectra, the measured contrast between the core of a spectral line and the continuum will be dependent on the temperature difference between the two regions in the planet's atmosphere. If temperature decreases with altitude (non-inverted atmospheres), the line cores will be formed at lower temperature than the continuum and will appear darker, producing *absorption* lines. On the other hand, if temperature increases with altitude, for instance due to a particular species absorbing incoming stellar radiation, the line cores will be formed at higher temperature than the continuum and will appear brighter, producing *emission* lines.

Not only is cross correlation sensitive to the shape and shift of the spectrum, but also to its sign. If we were to cross correlate a spectrum containing emission lines with an identical spectrum containing absorption lines (e.g. by changing the prescriptions for how temperature changes with altitude in the model), we would obtain almost perfect anti-correlation. It is thus straightforward to determine if thermal inversion layers are present from HRS data, and this was indeed recently done to determine the presence of titanium dioxide causing an inversion layer in the atmosphere of WASP-33 b [61].

Note that we stated “almost perfect” anti-correlation. In reality flipping the sign of the lapse rate, which is defined as $dT/d\log(p)$, does not correspond to an exact flipping of the output spectrum. Furthermore, the exact shape of a spectral line reflects the possibly complex behaviour of the lapse rate. Moving away from the line core to the line wings, it is possible to probe progressively deeper in the planet's atmosphere, therefore the line profile will become potentially very complex if e.g. the atmosphere is non inverted up to a certain altitude and inverted higher up. A schematic representation of the effects of various lapse rates on the output spectra and cross correlation functions is shown in Figure 59.

The complexity of line shape was one of the possible causes for the non-detection of CO in the emission spectrum of HD 209458 b [80]. While subsequent work has recovered

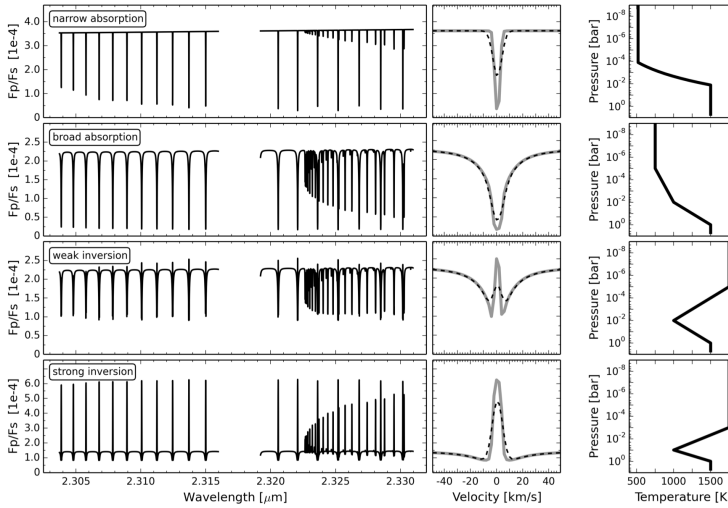


Figure 59: The complex dependence of the planet line profile on the temperature-pressure structure of exoplanet atmospheres, illustrated via model emission spectra of HD 209458 b ([80]). The left panels show the planet/star flux ratio as a function of wavelength, while the central panels show the resulting auto-correlation function, both at the original resolution of the models (grey line), and at the instrumental resolution of $R = 100,000$ (dashed lines). The corresponding temperature-pressure profiles are shown in the right panels, demonstrating the strong sensitivity of the line shape to the presence of inversion layers.

this detection [13], the signal in this planet remain incompatible with both an inverted and a non-inverted atmosphere with a strong lapse rate, and points instead to a milder lapse rate.

2.6 From demonstration to comparative exoplanetology

In Section 2.4 we mentioned that a sample of about 25 exoplanets has been investigated so far with HRS. We will now discuss whether the method can be applied on a much larger scale, thus opening up comparative exo-planetology in a way that is similar to what is currently possible with the dozens of transiting exoplanets observed with HST.

The first element to consider is the increase in instrumental capabilities that happened in the last few years. In reference to the signal-to-noise formula for the cross-correlation method, Equation 49, there are three main areas with direct impact on the sensitivity of HRS:

- **Telescope size:** as the number of collected photons (N_γ and $N_{p,\gamma}$ in Equation 49) grows linearly with telescope area, it follows trivially that the bigger the mirror, the higher the S/N. There is an additional dependence on the instrumental *efficiency*, i.e. the fraction of photons recorded by an instrument with respect to those incident on the telescope mirror. Modern spectrographs can achieve efficiencies of 5-10% in the near infrared, i.e. 2-3 times higher than the previous generation used for most of the HRS studies;
- **Resolving power:** the exoplanet signal ($N_{p,\gamma}$) has an additional dependence on the resolving power of the instrument. The reason for this is that spectral lines are broadened by the so-called instrument profile (IP) when observed through a spectrograph. Typically this profile can be approximated as a Gaussian profile with full-width half-maximum (FWHM) equal

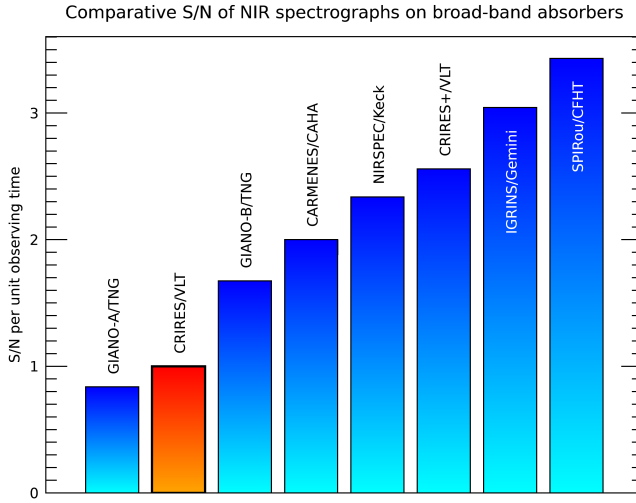


Figure 60: Comparison of the expected performances of past and existing near-infrared spectrographs (signal to noise per unit time) based on the criteria described in Section 2.6. CRIRES at the VLT is highlighted in orange as it was used to produce the very first successful detections.

to the resolution of the instrument. For the majority of exoplanets, excluding possible effects of winds and rotation (Section 3.5), the IP is broader than the width of the planet line profile. This holds up to $R \approx 250,000$ for hot Jupiters, and up to $R \approx 500,000$ for terrestrial exoplanets [48]. By comparing these values with Table 3, it is clear that the IPs of most of the spectrographs currently in operation will dominate the width of the observed profile. In this regime, the equivalent width of a spectral line is inversely proportional to the resolving power, thus the peak-to-continuum line contrast grows linearly with R . For higher values of R the planet line profile starts to be resolved, and therefore there is no further gain to increase the resolution;

- *Spectral range*: Equation 49 contains a term dependent on the number of spectral lines available for cross correlation ($\sqrt{n}$). For molecules that are opaque over a wide wavelength range, such as H_2O , CH_4 , and many of the species with 3 atoms or more, n scales linearly with the spectral range of the instrument. This is expected: the more simultaneous coverage a spectrograph has, the higher the number of lines that it can observe at once, thus increasing the gain in signal coming from the application of the cross correlation.

Figure 60 puts all this information together and ranks existing instruments in terms of their expected signal-to-noise ratio in cross correlation per unit time spent on sky. Due again to Poisson statistics, the S/N approximately scales with the square root of the observing time, thus a spectrograph with twice the score in the Figure needs $4\times$ less observing time to achieve the same detections. It is clear that telescope size is not the primary requirement of HRS, as past observations with CRIRES at the VLT ($R = 100,000$, 8.2-m mirror, about 2% throughput, 70 nm spectral range) can be outmatched by instruments mounted at smaller telescopes (e.g. 3.5-m class telescopes such as CAHA or CFHT) but covering the whole NIR (SPIRou) or a large fraction of it (CARMENES) with relatively high efficiencies (5-10%). The commu-

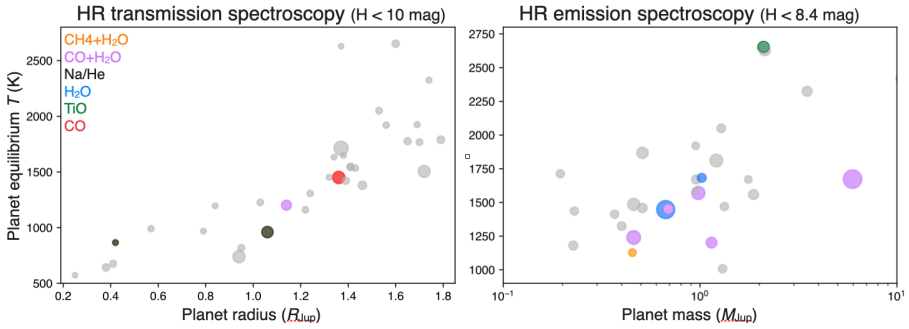


Figure 61: Known exoplanets for which atmospheric characterisation via HRS is well within the range of current instrumentation, i.e. a detection can be achieved in 5-10 hours of telescope time. Indicated in colours are those planets for which species have been detected already. Planets are selected for transmission spectroscopy (radius-temperature plot, left panel) and for emission spectroscopy (mass-temperature, right panel).

nity is just starting to tap into the power of medium-size observatories with HRS capabilities [3, 12, 34, 79].

The favourable performances of current instrumentation reflect into a much broader sample of exoplanets that can be targeted through HRS observations. By looking at Figure 61, it is clear that only a small fraction of the 30-40 exoplanets within reach of current observing facilities has been observed. While the detected sample is mostly focused on hot Jupiters or Saturn-size planets, the potential sample ranges from warm Neptunes to ultra-hot Jupiters, offering the unique opportunity to explore exoplanets with a wide range of physical and chemical conditions.

2.7 Follow-up of TESS exoplanets

Launched in April 2018, the Transiting Exoplanet Satellite Survey (TESS) mission [73] was designed at a relatively low-cost to find transiting exoplanets around the brightest stars. The mission is currently ongoing and has secured funding to extend its operation beyond the original two years of nominal length. At the time of writing, the satellite has nearly completed its primary mission to observe both the southern and northern ecliptic hemispheres, with over 1,900 candidates found and about 50 confirmed through follow-up observations. The mission is particularly suited to find planets of intermediate size ($2\text{--}4 R_{\oplus}$), which as we stressed already are the most common exoplanets [30]. Furthermore, a big fraction of these planets will be temperate and potentially in the classic Habitable Zone, defined in Section 5.1.

We can perform a simple exercise trying to scale the current sensitivity of HRS observations to the predicted yield of the TESS mission [95]. If we assume observations of water vapour in transmission and cloud-free atmospheres, the atmospheric signals can be easily scaled by using the formalism explained in Section 1.5.1 and accounting for planet equilibrium temperature, gravity, and increase in mean molecular weight. Further information is added about instrumental and telescope performances, following the considerations of Section 2.6, and a full year of dedicated observations is assumed, i.e. the planet is observed every time there is a transit opportunity with a penalty of 50% due to weather losses. Figure 62 shows the outcome of this simulated observing campaign for two fixed locations on the Earth (Mauna Kea, Hawaii and Cerro Paranal, Chile), and two state-of-the-art near-infrared spectrographs (SPIRou at CFHT and CRIRES+ at VLT).

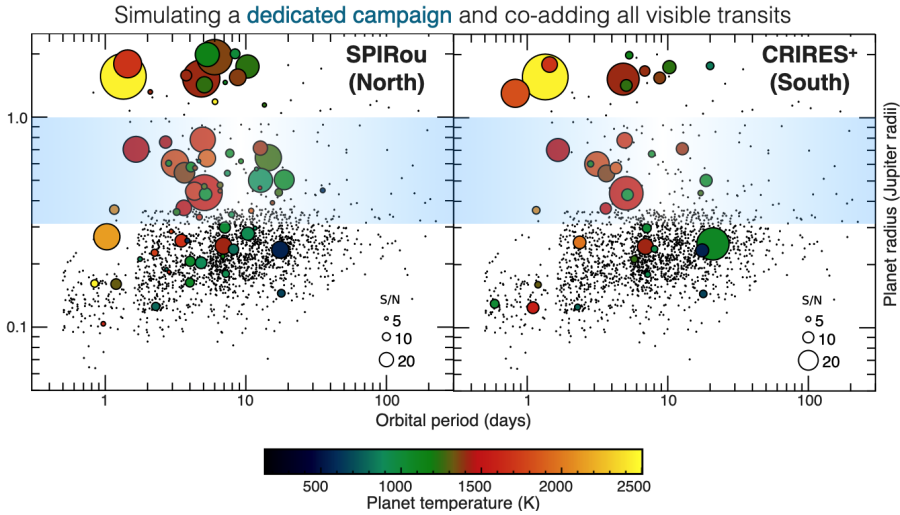


Figure 62: Simulated yield of the TESS mission highlighting the exoplanets that can be followed up via high-resolution transmission spectroscopy in the infrared (coloured circles). The S/N of the detection is shown by the size of the symbol, while its colour encodes planet temperature. While most of the giant exoplanets (above the band shaded in light blue) will be known already, there is a well populated region of planets between 3 and 10 $R_{\oplus}$ that can be followed up with a dedicated observing season. Simulations are shown for two observatories on the Earth, CHFT/SPIRou in the Northern hemisphere (left panel) and VLT/CRIRES+ in the Southern hemisphere (right panel).

The striking result from these simulations is that there will be a few dozens TESS exoplanets that can be characterised via HRS. In fact, this potential population is so big that unless we dedicate an entire facility to it (with a time investment of 500-700 hours per season), it is unlikely to follow up all of them. While the biggest of these simulated exoplanets will be known already from previous surveys (and will therefore be some of the targets in Figure 61), the majority of smaller exoplanets down to 3 $R_{\oplus}$ will be previously unknown. Furthermore, it appears that there is a handful of objects with radii between 1 and 2 $R_{\oplus}$ within the reach of a survey, albeit at low S/N. It is important to realise that these numbers are based on a simulated yield, and therefore such objects could be the fortuitous outcome of a small planet orbiting a particularly nearby, cool star (see Section 5.1) and such planets might not exist in reality.

In conclusion, between the existing sample and the exoplanets found by TESS the next few years will likely see an explosion of HRS observations, bringing the amount of detected atmosphere on the ballpark of 50-100 by the turn of the decade.

2.8 Cloudy atmospheres at high spectral resolution

The importance of studying the properties of Neptune-size exoplanets has been stressed already in the introduction of this Chapter (Section 1.1). We have stressed how aerosols (clouds or hazes) seem to be an ubiquitous feature of their spectra, and hinted to the fact that they can complicate their interpretation. In Section 1.5.1 we have gone more into details and explicitly mentioned the “muting” effect of clouds on the transmission spectrum of exoplanets. From

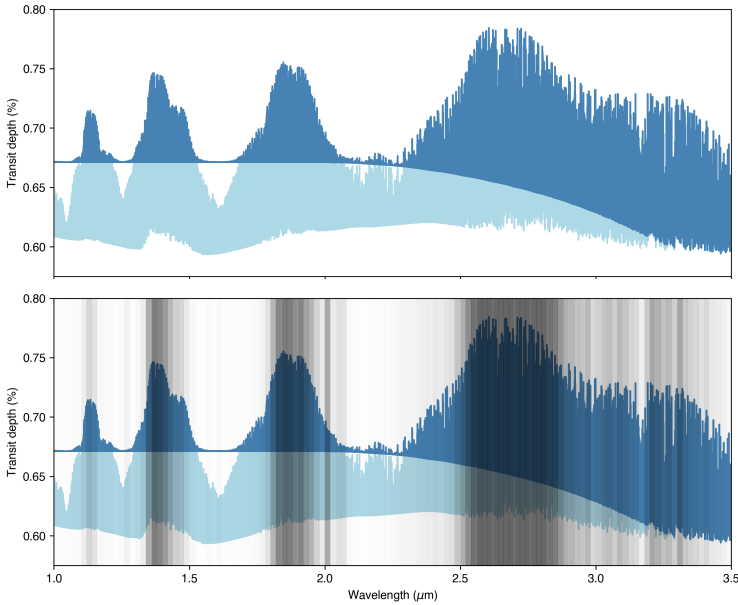


Figure 63: Top panel: Infrared high-resolution transmission spectrum of exoplanet GJ 3470 b in a cloud-free scenario (light blue) and with cloud parameters taken from literature observations (darker blue). It shows that clouds have a clear muting effect on the spectral features, although a significant fraction of the signal survives at the resolution of the model ($R = 50,000$). Bottom panel: same as the top panel, but with the transmission of the Earth’s atmosphere indicated through vertical opaque bands (darker = less transparent). It is clear that, due to water vapour in the Earth’s atmosphere, the best exoplanet signal is shielded from a ground-based observer.

Hubble Space Telescope observations, we have seen two extreme examples of this, i.e. the flat transmission spectrum of the super-Earth GJ 1214 b [46], and GJ 436 b [41]. More recently, we have also seen a less extreme case where the spectrum is heavily muted, but some of the molecular signature of water vapour survives [6].

In this Section we will investigate whether cloudy spectra look different at high spectral resolution, and we will highlight some important observational consequences. The top panel in Figure 63 shows two model spectra of GJ 3470 b, a Neptune size exoplanet with $T \sim 800$ K. In lighter colours we show the spectrum in the absence of clouds, whereas in darker colours the same spectrum is calculated with a cloud deck compatible with the observations of Benneke et al. [6], including a drop in the cloud opacity for wavelengths longer than 2.5 micron. This is expected when the size of the condensate particles becomes comparable to the wavelength of the light, because in this regime Mie scattering operates, which has an inverse dependence on wavelength.

Compared to the HST spectra, which is already almost completely flat at a resolving powers of $R = 35\text{--}50$, the HRS spectrum preserves a large fraction of the water vapour signal above the cloud deck, i.e. the signal of the strong spectral lines formed at atmospheric pressures lower than 0.1–1 mbar, which is the value inferred from HST observations. To be slightly more qualitative, while the amplitude of the water band around $1.4\text{ }\mu\text{m}$ is reduced by

about 90% in HST observations, simulated HRS observations show instead that more than 50% of the signal survives.

While the above finding seems in strong favour of high-resolution spectroscopy, it is important to realise that the surviving signal is just one part of the puzzle, because it neglects the fact that HRS can only be performed with ground-based telescopes. The bottom panel of Figure 63 shows the same two spectra of GJ 3470 b and in addition an indication of the main absorption bands in the Earth's atmosphere. Since water is one of the main constituents, the best observing windows are all outside of the main water bands (the classic *YJHK* infrared bands), but these are also the bands most strongly affected by clouds, even at high spectral resolution.

HRS allows us to partially overcome the blocking effects of the Earth's atmosphere and access portions of the spectrum traditionally excluded from broad-band photometry. This happens because the high resolution telluric spectrum has enough room in-between the strong telluric lines to allow astronomical observations. However, telluric contamination will need to be dealt with and the analysis described in Section 2.3 pushed much further than in the past. A better alternative is perhaps to aim for the spectrum at longer wavelengths, e.g. the *L*- and *M*-band accessible with CRIRES+ at the VLT or NIRSPEC at Keck. As we already mentioned, at these wavelengths the cloud opacity is likely to drop due to the effects of Mie scattering.

The verdict on whether high-resolution spectroscopy has a clear advantage over low-resolution observations is still pending, and while at the time of writing there are two studies in preparation showing that this is indeed the case (S. Gandhi, C. Hood, private communication), it is still premature to report these finding in the context of this Chapter.

3 The star as a source of astrophysical noise

Stars are far from being perfect black bodies. Even the earliest types - which at optical wavelengths show the smoothest spectrum, have spectra composed by an “envelope” resembling a black-body spectrum, which we call the continuum, and series of spectral lines. Lines in stellar spectra are formed by atoms or molecules present in the atmosphere, and their strength and richness are strongly dependent on the chemical composition and physical conditions of the star.

In previous sections we have already mentioned that moving in and out of a spectral line means changing the opacity of the atmosphere, therefore we expect different lines to be formed at different depths, each characterised by a certain temperature and a certain velocity. The latter is due not only to convective motion of stellar photospheres, but also to the component of the rotational velocity projected along the line of sight. It follows that stellar spectra encode the combined effects of chemical, physical, and dynamical conditions some of which can be variable on both temporal and spatial timescales. We should consequently expect that stellar spectra influence our inference on exoplanet spectra, which are measured using spectral and temporal differential measurements, especially during transit when an exoplanet crosses different portions of a stellar disk.

In HRS observations, a particularly relevant case is when studying a molecule, such as carbon monoxide, that has a very similar spectrum regardless of temperature. For instance, Figure 64 shows that the CO spectrum of an exoplanet at 1,800 K (bottom panel) is quite similar to the CO spectrum formed in the photosphere of G or K dwarfs (top panel). This means that any residual stellar spectrum not effectively removed by the analysis explained in Section 2.3 can result in spurious cross correlation signals competing with, and in some cases outshining, the planet signal. We therefore expect stellar spectra to produce a complex Doppler signature and to be a “nuisance” to exoplanet observations.

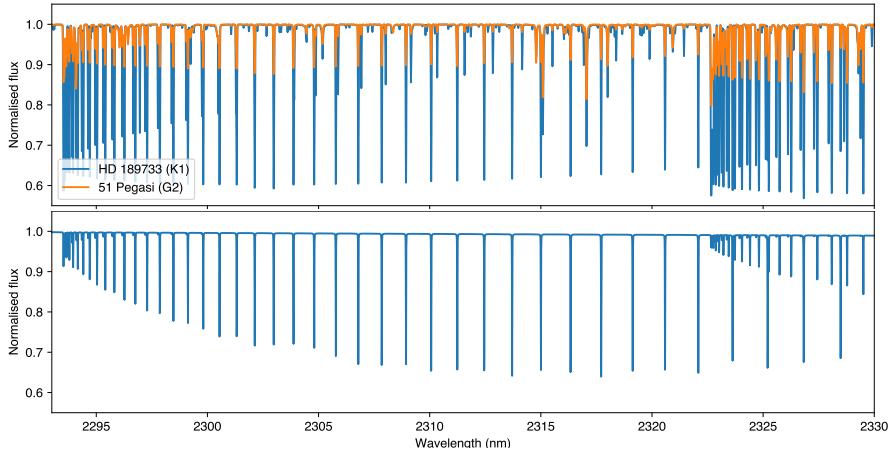


Figure 64: Spectral signature of carbon monoxide around 2.3 micron. The top panel shows two main sequence, solar abundance stars (the G2 dwarf 51 Pegasi and the K1 dwarf HD 189733), whereas the bottom panel shows the emission spectrum of a hot Jupiter at 1,800 K. Despite the big difference in temperature, the structure of the CO band is remarkably similar.

In the next sections we will discuss the various effects that can produce a net stellar Doppler signature interfering with exoplanet signals.

3.1 Stellar spectra are non-stationary

In Section 2.3 we have shown how the analysis applied to high resolution spectra can effectively remove any spectral component that is stationary in wavelength, i.e. with a negligible change in radial velocity during the observations. The issue here is that the star will change its radial velocity slightly during a few hours of observations, due to two main effects:

1. The reflex motion around the centre of mass of the system hosting an exoplanet. For hot Jupiters, this can amount to a few tens of m s^{-1} and is used to detect exoplanets via the radial velocity technique (Section 1.3). Due to the fact that HRS of exoplanets deals with Doppler shifts at the level of km s^{-1} or greater, the stellar RV component is generally regarded as negligible;
2. The change in *barycentric* velocity of an observer on the Earth. The latter will move with a non-zero radial velocity with respect to the centre of gravity of the solar system, which is called barycentric velocity (v_{bary}) and is time-variable. It will add up on top of the radial velocity of the system (the systemic velocity v_{sys}). The barycentric velocity consists of the projected rotational velocity of the Earth (at most $\pm 0.5 \text{ km s}^{-1}$ at the equator) and the projected orbital velocity of the Earth (at most $\pm 30 \text{ km s}^{-1}$). The combined change of these two components over a few hours of observations typically amounts to a few hundreds m s^{-1} .

The combined radial velocity of the star measured by an observer on the Earth ($\text{RV}_\star$) is thus

$$\text{RV}_\star = v_{\text{sys}} + v_{\text{bary}}(t) + K_\star \sin\{2\pi[\varphi(t) + 0.5]\}, \quad (53)$$

where the negligible orbital component is the product of the stellar RV semi-amplitude ($K_\star$) and the sine of the orbital phase of the star. Stellar phases have been obtained from the planet's orbital phase φ by adding 0.5, which is half of the orbit.

We can now write a complete version of Equation 32 expressing the planet's radial velocity as measured by an observer on the Earth:

$$RV_P = v_{\text{sys}} + v_{\text{bary}}(t) + K_P \sin[2\pi\varphi(t)], \quad (54)$$

where K_P is now the semi-amplitude of the planet's radial velocity curve. In the two equations above we have explicitly indicated those terms that are time dependent.

3.2 Stellar cross-correlation noise in day-side observations of exoplanets

Due to the v_{bary} term in Equation 53, the stellar spectrum is non-stationary in the reference frame of a ground-based observatory, and it will not be perfectly corrected by the analysis applied to HRS data. Imperfect correction will result in residual cross correlation noise, which is evident in the two-dimensional cross correlation map shown in Figure 65 (top panel). This is a real observation of CO in the atmosphere of the exoplanet HD 189733 b, originally published by [24] and re-analysed by [21]. In theory, due to the fact that the star moves with the system but its RV amplitude is small ($K_\star \ll 1 \text{ km s}^{-1}$), the stellar noise should be confined to $(v_{\text{rest}}, K_P) \approx (0, 0) \text{ km s}^{-1}$. However, due to the change in RV caused by the v_{bary} term and by the fact that even small stellar residuals can be orders of magnitude stronger than the planet's signal, spurious stellar cross correlation propagates to $K_P \neq 0$, causing an evident nuisance in the data even at the position of the planet (white dashed lines in the Figure). In this case the detection of the planet would not be prevented, but certainly confused by the stellar noise.

A more serious case is the study of the emission spectrum of exoplanet 51 Pegasi b, originally published by [14] and also re-analysed by [21]. Being a non-transiting planet, the orbital radial velocity of 51 Peg b is not known. When co-adding three different nights of observations, there is a candidate exoplanet signal at the expected null rest-frame velocity and $K_P \approx 130 \text{ km s}^{-1}$, however this is also strongly contaminated by stellar residuals in one of the nights (Figure 66, left panels). This contamination prevented the authors of the original paper from claiming a secure detection.

In Section 3.4 we will show how these observations can be cleaned up by using a three-dimensional model of the stellar photosphere to accurately remove the stellar lines.

3.3 Stellar spectra are distorted during transit

The noise from the underlying stellar spectrum is even more severe in the case of a transiting planet. This is due to the fact that during transit the planet is blocking different parts of a rotating stellar surface, and that this surface is heterogeneous in that the intensity of the spectrum varies from the centre to the limb. Lastly, the stellar photosphere is subject to granulation and activity, which further breaks any assumption of homogeneity. In this Section we will treat each of these effects separately.

Moving from the centre to the limb of a star, we see progressively higher layers of the stellar atmosphere. This is due to a geometric effect: while the photosphere of the star emerges from the layer with optical depth close to unity, due to the slanted angle between the normal to the stellar surface and the line of sight the same distance along the line of sight will correspond to a smaller depth into the stellar atmosphere to keep the optical depth constant. As a consequence, the layers from which the stellar radiation emerges will be cooler at the

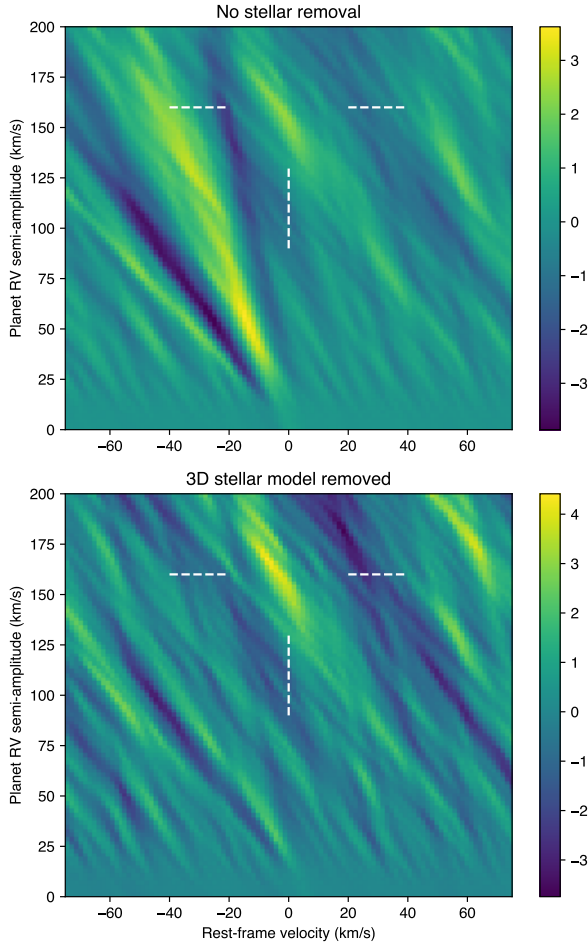


Figure 65: Top panel: Detection of CO absorption in the dayside spectrum of exoplanet HD 189733 b (de Kok et al. 2013) as a function of planet rest-frame velocity and radial-velocity semi-amplitude. While the planet is detected at the expected position (white dashed lines), it is also contaminated by strong stellar residuals, as shown by the slanted broad features blue-shifted compared to the planet’s signal. Bottom panel: same as the top panel, but with the stellar spectrum now removed through the modelling described in Section 3.4.

limb (resulting in the typical darkening quite evident at optical wavelengths) while spectral lines may result deeper or shallower according to their preferential formation depth. We generically refer to this effect as Centre-to-Limb Variations (CLVs) and they can represent a significant source of excess (or defect) flux in the transit light curves, at the level of precision compatible with atmospheric studies [18].

A rotating stellar surface produces a broadening of the stellar spectrum, with spectral lines deviating from the quasi-Gaussian profile that would be produced by the Instrument Profile, and converging instead towards a bell-shaped profile for progressively faster rotation. Since each portion of the rotating disk contributes to the spectrum with a different Doppler-shift (a different portion of the line profile), when an exoplanet blocks a fraction of the stellar disk

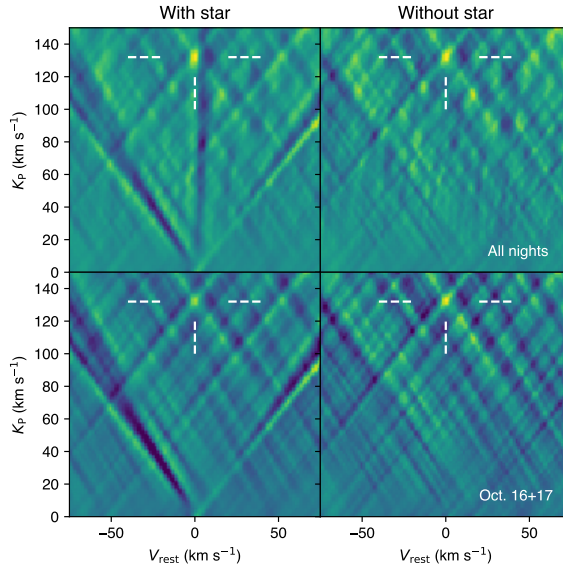


Figure 66: CO absorption in the dayside spectrum of exoplanet 51 Pegasi b as a function of planet rest-frame velocity and radial-velocity semi-amplitude, originally published by [14] as a tentative detection. They recovered a signal similar to that shown by the left panels, where in addition to a well localised planet candidate (white dashed lines) there was a noticeable stellar residual signal (the slanted lines at different angles). Once a model of the stellar spectrum has been removed as explained in Section 3.4 (right panels), the planet signal remains the only unambiguous detected source.

this profile results “distorted”, because it will be missing the contribution from the underlying fraction of the rotating stellar surface. This is known as the Rossiter-McLaughlin effect and is known since a century to affect the spectra of binary eclipsing stars [56, 78]. It has been used widely to measure the alignment between the spin of the star and the orbit of exoplanets [69], and it results in a spurious shift of the centroids of cross correlation functions in data taken for stellar radial velocities.

3.4 Modelling the background stellar spectrum

Stars more massive than 1.3 solar masses transport heat in their outer layers by convection. This can be simulated by three-dimensional models of stellar photospheres, such as the snapshot shown in the bottom panel of Figure 67 [53]. The red and blue stars denote elements on the stellar surface (columns of fluid) that rise toward the surface or sink back toward the interior, respectively. The corresponding profiles of a CO line formed at those two positions are shown in the top panel with the same colour scheme, and compared to all the other profiles in the snapshot. While on average the amount of mass rising and sinking need to conserve, the weighting of each parcel of fluid is not equal. Rising columns of gas are hotter (hence brighter) and they occupy more volume than the sinking columns, therefore they weigh more in terms of emitted flux. This means that overall we expect the profile of a spectral line formed in the photosphere to be blue-shifted by convection. The actual line profile is even

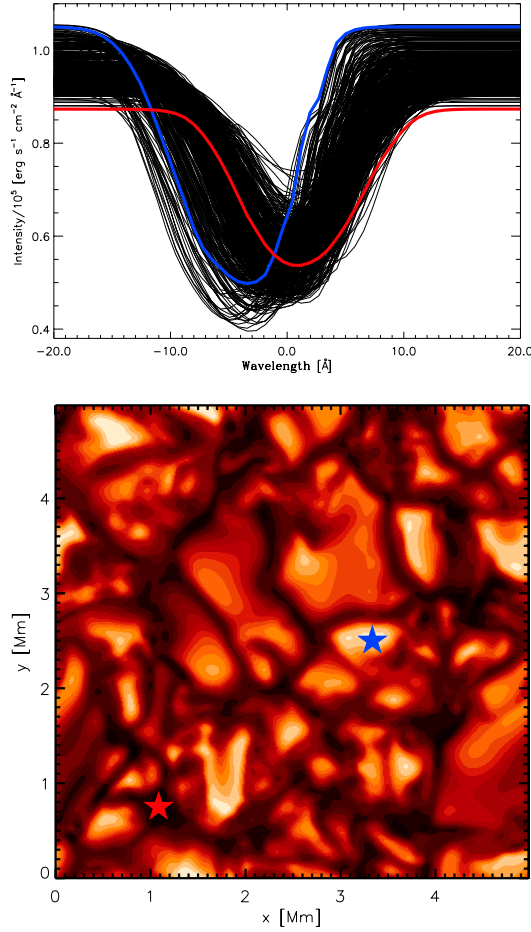


Figure 67: Snapshot of a three-dimensional K-dwarf simulation computed at the centre of the stellar disk, showing the granulation pattern of the star (bright and dark cells in the bottom panel). The varying profiles of the CO line around 2301.5 nm are shown in the top panel, with those coming from a rising (in blue) and a sinking (in red) column of fluid highlighted in colour. The observed CO profile, resulting from the sum of all the profiles, will be skewed and blue-shifted overall.

more complicated, and results from the addition of all the columns of fluid in the convective stellar photosphere, i.e. from the addition of all the profiles shown in the top panel of Figure 67. It follows that a realistic stellar profile, which can be appropriately encapsulated via 3D simulations, will be skewed and blue shifted overall. The more common practice to approximate a stellar line profile by using the outcome of one-dimensional radiative-transfer calculations, broadened by a micro- and macro-turbulence parameters, can only approximate the symmetric part of the line profile, but cannot account for its skewness and overall blue shift.

It is possible to extend the 3D simulations to the whole stellar surface by tiling this up with many boxes such as those in Figure 67, appropriately tilted in order to follow the direction normal to the stellar surface. Tilting a three-dimensional box will ensure that radiative

transfer is computed along slanted line of sights, thus accurately simulating the effects of CLVs. This will give us the stellar spectrum as a function of position on the stellar surface, i.e. a radial coordinate μ (the cosine between the normal to the stellar surface and the line of sight) and the azimuth ϕ . Once integrated numerically over the whole stellar surface, the modelled spectrum is ready to be removed from the observational data, provided the spectra have been accurately calibrated in wavelength and that the instrumental broadening profile of the spectrograph is well known.

When such a procedure is applied to the spectra of 51 Pegasi and HD 189733 previously discussed, the residual cross correlation coming from the stellar spectrum is reduced to negligible levels and it is possible to securely identify the planetary signature as the only significant source in velocity space. This is shown by the bottom panel of Figure 65 for HD 189733 b and by the right panels of Figure 66 for 51 Pegasi b.

Moving onto transiting planets, the model needs to be expanded to solve for the geometry of the transit, i.e. determining which elements of the stellar surface are occluded as a function of time. This can be done with various degrees of complexity, but here we review the simple numerical approach where we subdivide the stellar surface into a Cartesian grid, and we determine which elements are eclipsed by the planet at each instant of the transit. As every element of the grid has a spatially-resolved spectrum (resulting from the 3D simulations introduced above) and a velocity field (from stellar rotation) associated, the total spectrum is obtained by summing over all the unocculted elements, properly Doppler shifted to account for their radial velocity.

The power of this approach is that the simulated spectrum will naturally include the RM effect and CLVs, as shown for example in Figure 68 for HD 189733, a K1V dwarf, computed around the absorption band of CO while the hosted exoplanet HD189733 b transits. In the figure we have applied the principle of Doppler tomography, where the average spectrum out of transit is subtracted out of the entire sequence, thus revealing the changes in the line shape as a function of time as an excess or deficit of flux. It shows that for this system CLVs account for a residual signal at $\sim 5 \times 10^{-4}$ level, while the combined CLVs + RM effect amount to an order of magnitude more, or about 5×10^{-4} . This is also the order of magnitude of the signal expected from the planet's transmission spectrum, and therefore it follows that if RM and CLVs are not properly modelled and removed, no signal can be confidently detected from the planet. In the next section we will review a practical application of the modelling described in this section, that is the measurement of winds and rotation of an exoplanet from its transmission spectrum.

3.5 Stellar modelling to accurately measure exoplanet winds and rotation

Due to their close-in orbits, hot Jupiters become tidally locked on very short timescales. As shown by [84], the synchronisation timescale τ_{sync} can be approximated as

$$\tau_{\text{sync}} \approx Q \left(\frac{R_p^3}{GM_p} \right) \left(\frac{M_p}{M_\star} \right)^2 \left(\frac{a}{R_p} \right)^6 (\omega - \omega_s), \quad (55)$$

where Q is the quality factor describing how the planet's interior reacts to torques, a the semi-major axis of the orbit, ω the initial rotational frequency and ω_s its synchronous value. For typical hot-Jupiter systems and $Q \approx 10^5$ (the accepted value for Jupiter), the timescale is as short as 10^6 years, negligible compared to the age of evolved systems (> 1 Gyr). Synchronisation implies that the rotational and orbital periods are the same. This produces two main regimes of atmospheric circulation in hot Jupiters:

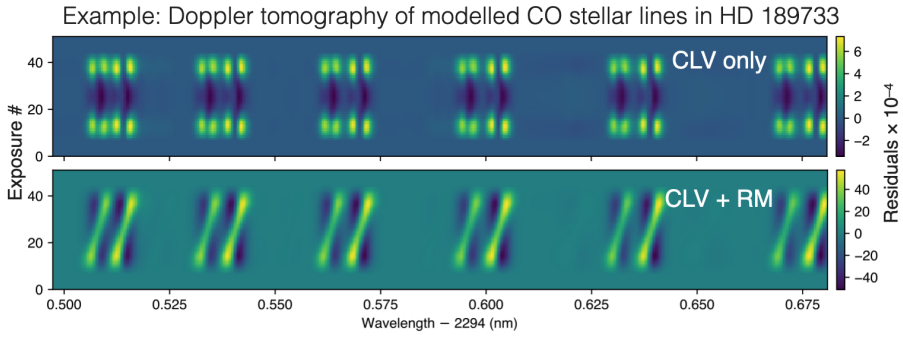


Figure 68: Modelled spectrum of HD 189733 (K1 dwarf) during the transit of its exoplanet, divided by the out of transit average model spectrum. The top panel shows the effect of centre-to-limb variations on the shape and intensity of CO lines, while the bottom panel combined this effect with the (much stronger) distortion coming from the differential occultation of the rotating stellar surface during transit (Rossiter-McLaughlin effect).

- At low pressures (higher up in the atmosphere), winds up to a few km s^{-1} flow from the permanent dayside to the permanent night side of the planet, resulting in a net blue-shift measurable during transit, i.e. when we probe the planet’s terminator;
- At high pressures (lower in the atmosphere), i.e. the 0.1-1 bar typical of the planet’s photosphere, the Coriolis force causes the onset of a strong eastward equatorial jet, which can be super-rotating compared to the bulk synchronous rotation. This causes a net shift of the planet’s maximum thermal emission towards the east, which observationally results in phase curves having maxima before secondary eclipses.

Interestingly for HRS, the interplay of atmospheric winds with intensities up to a few km s^{-1} can broaden and more generally distort the planet line profile, to an extent that is likely detectable through observations [40, 58, 71, 85]. In the optical the effect of winds has indeed been measured by [99] and [49] through the investigation of sodium absorption in the atmosphere of HD 189733 b. A few years earlier, [91] tentatively claimed a net blue-shift of the transmission spectrum of another archetype hot Jupiter, HD 209458 b. In this case, rather than looking at the sodium doublet, the blue shift was measured through the combined cross-correlation of all the CO lines around $2.3 \mu\text{m}$.

In this context we will briefly review the first attempt at constraining both winds *and* rotation from the transmission spectrum of a very well known hot Jupiter, HD 189733 b. This system has come up a few times before, and especially in Sections 3.1 and 3.4 when mentioning the necessity for modelling and removing the spectral component from the parent star.

A very first attempt at stellar removal was made by using a parametric line profile via micro- and macro-turbulence [10]. Due to the asymmetries in the line profile described in Section 3.4, this approach was not sufficient to completely correct for the stellar spectrum, however it was at least successful at suppressing the CO signal from the parent star, thus revealing a clear and unambiguous detection of combined H_2O and CO absorption coming from the planet’s atmosphere. Additionally, the analysis of the shape of the planet line profile, inferred from the shape of the cross correlation function, allowed the authors to claim a net blue shift of the exoplanet spectrum and a broadening most consistent with an equatorial wind speed around 3.5 km s^{-1} .

All this inference was obtained via a two-parameter model of the planet’s atmospheric circulation. The planet was supposed to rotate as a rigid body (yielding the first parameter — rotational period) and an additional equatorial super-rotation was allowed (yielding the second parameter — equatorial super-rotation). Due to the intrinsic similarity of these two parameters on the planet’s line profile, it was not possible to disentangle the rotation component from the equatorial jet super-rotating, but nevertheless it was possible to exclude fast rotations that would be more typical of the giant planets of the solar system.

However, it is important to realise that a parametrisation such as the one just described is not physically plausible. Since rotation triggers the onset of the Coriolis force, which in turns determine the global circulation patterns, winds and rotation are intrinsically connected and cannot be considered as independent. [28] took one step forward and utilised the predictions from three-dimensional general circulation models (GCMs) to calculate models to cross correlate with the existing data. As the diagnostic from GCMs adds a significant level of detail to the line shape, they also decided to pair the analysis with the correction of the stellar spectrum via the 3D modelling outlined in Section 3.4.

The resulting improvement in the measured signal from carbon monoxide alone is shown by the right panel in Figure 69, which also contains for comparison the same signal measured without any stellar correction and with the one-dimensional approach of [10]. Such a refined approach allowed the authors to improve the constraints from their parametric model of winds and rotation. Furthermore, it firmly showed that GCM models can correctly reproduce a net blue shift of the spectral lines during transit (due to the low-pressure atmospheric flow), and that indeed rotation and equatorial jets are physically inter-connected. As soon as rotation is significant, the equatorial jet stream onsets with wind speeds nearly independent from the rotational rate. This has the effect to “equalise” the width of the line profile, thus hindering our ability to precisely measure the bulk planet rotation. This result clearly shows the shortcomings of a purely parametric approach if the choice of parameters is not informed by a thorough physical modelling.

4 Estimating significance and error bars

In previous sections we presented a very qualitative review of HRS results, focussing on the analysis and on the scientific questions, without delving too much into the details of how significance and error bars were obtained. As this is not a trivial issue, this Section is devoted to clarify some of the techniques used in past literature.

At low spectral resolution the analysis produces spectra that can be directly compared to models, e.g. via chi-square analysis. This also enables relatively straightforward Bayesian retrieval via e.g. MonteCarlo simulations. HRS returns instead a matrix of cross correlation values, and these are notoriously harder to treat because we do not have a “ground truth” to take as a reference. Indeed, as we explained in Section 2.2, the actual observed spectrum is too noisy to show any spectral features, so we have to rely on measuring the level of correlation between the data (whatever spectrum they contain) and a set of models. Understanding the statistical properties of cross correlation functions is also an additional challenge, but thankfully all these aspects of the analysis are currently known, although much of the progress has only been made in the last few years.

4.1 The signal-to-noise approach

The first and most straightforward method to estimate the level of a detection relies on the *signal-to-noise* theory. In the two-dimensional matrix of CCF versus (v_{sys}, K_p) , such as those

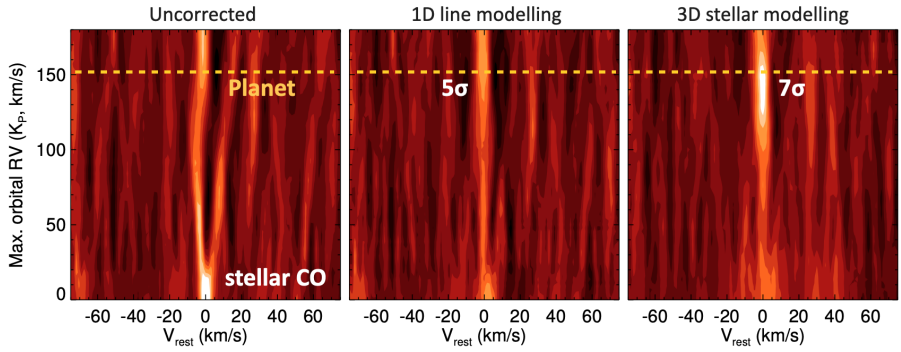


Figure 69: Carbon monoxide signal in the transmission spectrum of HD 189733 b as a function of rest frame velocity and planet maximum radial velocity, measured with three different analysis. In the left panel, the stellar spectrum is not modelled nor removed prior to the analysis described in Section 2.3, resulting in strong stellar residuals in cross correlation due to RM effect and CLVs. In the middle panel, a one-dimensional line profile is assumed for the stellar CO lines and the modelled spectrum is removed from the data, resulting in noticeable improvement ([10]). In the right panel, a full 3D simulation of the stellar photosphere as described in Section 3.4 is utilised to clean the data from the stellar spectrum, resulting in the complete correction of stellar residuals ([28]).

shown in Figure 58, the maximum value of cross correlation is divided by the standard deviation of the values away from the peak to obtain the signal-to-noise ratio (S/N) of the detection. Error bars are then defined by the region in the parameter space where the S/N drops by one. The S/N is an intuitive quantity to compute, however it has two major drawbacks. Firstly, some of the structures in the CCF away from the peak are not noise, but rather secondary peaks (also known as aliases) of the real signal. Secondly, it is plausible that different values of (v_{sys}, K_P) have different noise properties, because moving in velocity means sampling regions of the spectra that can be affected by telluric or stellar residuals differently (recall our discussion about the slope and the shift of the CCF in Section 2.3). Lastly, by computing the S/N in the total cross-correlation matrix (i.e. after summing in time), it is not possible to identify cases where the detection comes only from a few spurious CCFs (i.e. individual spectra with anomalously high noise) rather than from the whole sequence. While normally this last condition can be excluded by visual inspection, arguably a more objective practice would be preferable.

4.2 The Welch t-test

A more statistical assessment of the significance of the detection can be obtained by analysing the values of cross correlation before they are summed in time, i.e. on the CCF as a function of RV and time shown in Figure 57. If a correlation signal from the planet is present, the values of the CCF corresponding to the planet’s radial velocity curve (the so-called *planet trail*) will have a higher mean than those away from the planet’s RV. In the figure this is shown by the slanted line with darker colour values following the planet’s trail.

The Welch t-test allows us to perform a statistical test on two sample distributions — the in-trail and out-of-trail distribution of CCF in our case — and reject the null hypothesis that the two are drawn from a parent distribution with the same mean. The version of the t-test

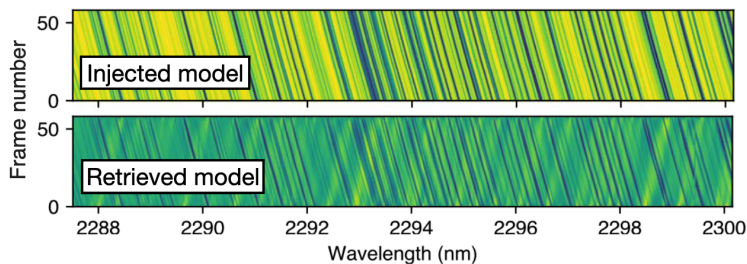


Figure 70: Example of the effects of HRS data analysis on the planet signal. A time sequence of exoplanet model spectra (top panel) is injected in real data at the early stages of the analysis, and then recovered at the end of the analysis (bottom panel) as explained in Section 4.3. Although largely preserved, the exoplanet signal has been visibly altered with extra structures appearing in both the temporal and wavelength domains.

designed by Welch is more general than the standard Student's *t*-test, in that it allows the two samples to have unequal size and variance.

Two assumptions are fundamental to the proper application of the *t*-test: the former is that the sample of cross correlation values has a Gaussian distribution, and the latter that the values are independent. While the agreement between the CCF values and a Gaussian distribution can be easily proven via Gaussian fitting to the histogram of cross correlation values, or more qualitatively via Q-Q plots, the assumption of independence is unlikely to be completely correct. Even by ensuring that the step in RV is coarse enough not to oversample the instrumental profile (which, for instance, is 3 km s^{-1} in the case of CRIRES), some level of correlation between adjacent radial velocity values is inevitable. In applying the *t*-test it is implicitly assumed that this level of correlation is small and most importantly constant, so that it affects all the data equally. This is a very good approximation due to the good stability of instrumental profiles.

The level of significance at which the null hypothesis is rejected (i.e. the in-trail and out-of-trail samples have the same mean) is quoted as the significance of the detection, and $n - \sigma$ error bars can be determined as contours levels in the parameter space where the significance drops by n .

4.3 Estimating the “model” cross correlation function

Some applications of HRS require not only matching the position and amplitude of the planet's spectral lines, but also their shape. This is necessary e.g. when estimating winds and rotation of exoplanets such as in Section 3.5. In this case a different approach was introduced to estimate the cross correlation of a model directly, so that it can then be compared to the measured cross correlation via goodness of fit statistics.

In the case of a perfect match between the planet spectrum and the model spectrum, one would be tempted to assume that the measured cross correlation should be equal to the model auto-correlation function (that is, the model cross-correlated with itself), except for the noise budget. However, due to the data analysis applied to remove telluric and stellar lines (see Section 2.3), the true planet spectrum is altered.

Fortunately, a fairly robust method can be used to quantify the exact nature of this alteration. A planet model spectrum, scaled to the right planet/star flux and shifted by the

appropriate radial velocity following Equation 54, is injected into the data as early as possible along the pipeline (normally just after wavelength calibration). By injecting here we mean multiply by the transmission spectrum in the case of transit observations, or by $(1+F_p/F_\star)$ in the case of an emission spectrum. The injected dataset is then passed through the full data analysis to remove telluric and stellar spectra, and lastly the difference between the injected and the real data is taken. As shown in Figure 70, while largely preserved by the analysis, the planet signal has been visibly affected, most notably in the form of extra banding or stretch as a function of wavelength and time.

We can take this comparison a few steps further, and take the difference between the cross correlation functions after summing the whole sequence in time. This will effectively give us the cross-correlation function of our model, which we can now compare to the CCF measured from the real data via chi square. This approach was used successfully in past work [10, 13, 28] and remained the most effective approach until the recent development of a full statistical treatment of the CCF.

4.4 Bayesian analysis on high-resolution spectra

The most recent approach at deriving the statistical properties of cross correlation in its exoplanet application was done by [15]. They took a more direct approach where they converted cross-correlation values into a likelihood function. In a “standard” goodness of fit approach where models are compared to data directly it can be shown that there is a direct relationship between the chi square and the log-likelihood, such as

$$\log L = -\frac{1}{2}\chi^2. \quad (56)$$

When data and model are cross correlated, a more complicated relation can be derived:

$$\log L = -\frac{N}{2} \log(s_f^2 + s_g^2 - 2R), \quad (57)$$

where s_f^2 and s_g^2 are the data and model variance, respectively, R is the cross covariance (i.e. the numerator of the cross correlation), and N is the length of the vector over which a cross-correlation function is calculated.

With a cross-correlation-to-likelihood conversion scheme, it is then possible to run Bayesian retrieval, e.g. MonteCarlo methods, to estimate the best-fit parameters and determine their confidence intervals. As conceptually simple as it sounds, the implementation took about two years of testing, and most of the work went into ensuring that the analysis would be unbiased. In a nutshell, the distortion effects on the planet signal described in Section 4.3 need to be minimised through a specific analysis, and replicated on the model as well.

In the optical, where the instrumental and atmospheric stability is superior, alternative likelihood conversions have been derived following [15], for instance a similar approach by [32] where the authors allow for unequal errors on each element of the vectors to be cross correlated, or [65], where the likelihood is reduced to a chi-square on the cross correlation functions.

In [15], strong conclusions were derived on the investigative power of HRS from simulated datasets:

- HRS can measure *absolute* abundances, and not only relative abundances. On known hot Jupiters, sub-dex precision on abundances is achievable.

- HRS can constrain radial velocities of currently observed exoplanets to better than 1 km s^{-1} . It is thus potentially sensible to the large-scale atmospheric circulation induced by rotation and radiation forcing on these exoplanets (see e.g. the measurements described in Section 3.5).

4.4.1 Combining low- and high-resolution spectroscopy

With a cross-correlation-to-likelihood scheme available, it becomes trivial to combine observations from space observatories (photometric points or low-resolution spectra) with HRS. As long as the same set of models is computed consistently for all data sets (i.e. with the same assumptions and numerical methods), it is just sufficient to sum the log-likelihood functions and compute Bayesian inference on the total log-likelihood function.

Combining datasets at varying resolution is not just a signal-to-noise exercise. Space observatories excel in delivering precise absolute measurements of the planet-to-star flux, and broad-band variations. High-resolution spectroscopy delivers information about the core-to-continuum line contrast, line shape, and line-to-line contrast. These are two complementary sets of information, i.e. they are two *independent* measurements encoding the same atmospheric parameters, thus leading to error bars that are shaped very differently. Therefore, when computing the joint distribution of these data, the expectation is that the error bars will shrink significantly not only because more data is added, but also because part of the parameter space is mutually excluded.

While [13] hinted at this synergy between low- and high spectral resolution even without the rigorous use of a likelihood function, it was in [15] that simulations clearly showed the advantage. The hot Jupiter HD 209458 b was simulated both via HST/WFC3 observations of secondary eclipse (at 30 ppm precision per 35 nm bin), and VLT/CRIRES observations of dayside spectrum (5 hours). By combining the datasets it was possible to constrain the planet's metallicity to within 1 dex and the C/O ratio to within 0.3 dex. The same exercise was repeated on real data by [31], who reported similar improvements in the precision of measured abundances and $T - p$ profile.

With this tool at hand, it is now possible to simulate coordinated ELT and JWST observations, to determine not only what is the optimal combination of exposure time and wavelengths to extract the maximum information from the data, but also what are the scientific questions that most benefit from the combined approach.

5 Bridging the gap towards habitable planets

5.1 Planets around M-dwarf stars

M-dwarf stars represent the low-mass tail of the stellar population. As such they are *smaller* and *cooler* than the Sun, and therefore much less luminous. While this generally results in stars that are quite faint at optical wavelengths, in the near-infrared M dwarfs — especially the closest ones — can still be quite bright.

The opportunity for exoplanet characterisation becomes clear if we consider the planet/star flux ratio of a hot Jupiter around a sun-like star, and we ask ourselves what type of exoplanets would be observable around an M dwarf by keeping the ratio constant. Figure 71 shows a schematic view of this exercise, from which it is possible to conclude that a 1.3% transit depth around a low-mass star would allow us to detect a planet that is just twice the size of the Earth. But the increase in sensitivity becomes even bigger when we consider emission spectra, as the thermal emission of a 1,200 K hot Jupiter around a solar type star would be equivalent to that of a planet with 4 times the Earth's radius and temperature equal to 800 K

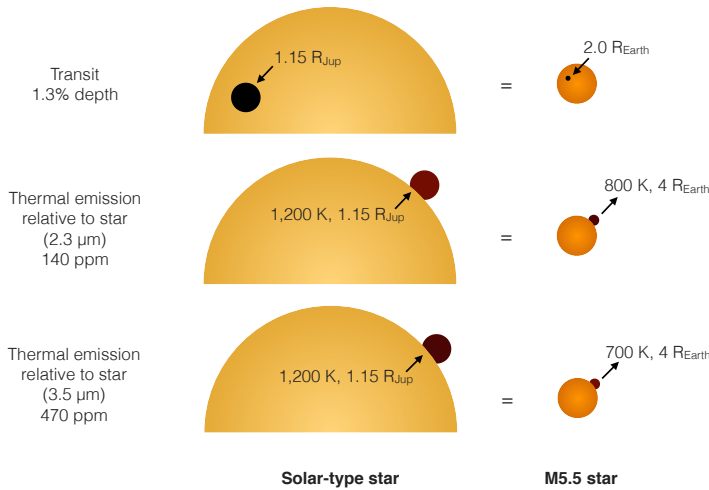


Figure 71: Schematic comparison of transmission (top row) and emission signals (middle and bottom row) from exoplanets, showing that a hot Jupiter around a solar-type star has approximately the same planet/star contrast than a warm Neptune around a M-5.5 star.

(at $2.3 \mu\text{m}$) or 700 K (at $3.5 \mu\text{m}$). This striking result tells us that warm sub-Neptunes around nearby M dwarfs are already within the reach of current technology.

The question is: are there such stars and such planets? The answer to the first question is definitely yes, as M dwarfs happen to be the most numerous stars in the solar neighbourhood, making up over 75% of the stellar population. To answer the second question, we can refer to the study of [25], who clearly highlighted that while exoplanets are common around any stellar type, their occurrence rate firmly increases around M-dwarf stars to yield around 2.5 planets per star. A quick census of nearby exoplanet systems seems to confirm these predictions: despite being a highly incomplete sample, there are 70 planets known within 10 pc, and 35 within 5 pc, with notable examples being the couple of exoplanets orbiting the nearest star to the solar system, the M5.5 Proxima Cen [4], and the system of 7 transiting planets around the M8 star Trappist-1 [33].

A third element conspires in favour of planets around M-dwarfs, which is their low luminosity. The range of distances from the parent star compatible with the long-term presence of liquid water on the surface of a planet identical to the Earth (i.e. the classic Habitable Zone — or HZ) moves inward with decreasing luminosity, with the consequence that temperate planets in the HZ of M dwarfs have very tight orbits [45]. For instance, while Proxima Cen b orbits at only 0.047 au from Proxima, it is firmly in the HZ. This has a very favourable observational consequence: not only are temperate planets around M dwarfs orders of magnitude brighter than around solar-type dwarfs when compared to their parent star, but they also have an increased probability of transiting, as the transit probability scales approximately inversely with the semi-major axis of the orbit. Transit observations can be repeated every few days (rather than once per year) due to the short orbital periods, quickly building up the signal to noise required to detect species in these exoplanets. All in all, populations studies suggest that 30% of M-dwarfs should host an exoplanet in the classic HZ [25]. It is therefore not surprising that the first temperate rocky planets that will be studied in the future to search for life will be around M-dwarf systems. In the remainder of this Chapter we will review the type of measurements that HRS will enable on planets in the classical HZ zone.

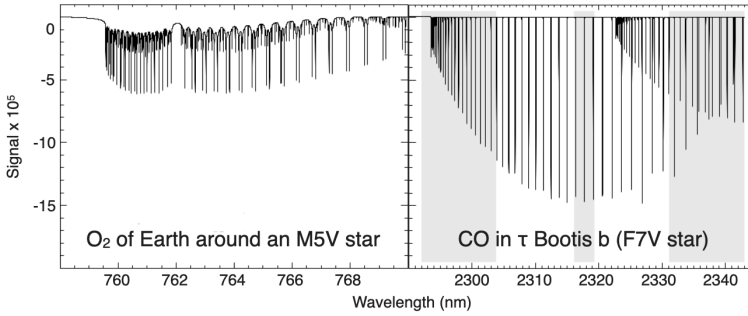


Figure 72: Simulated transmission spectrum of an Earth-like planet orbiting an M-5.5 dwarf star, around the oxygen A band (left panel). This is shown on the same scale as the CO absorption signal measured from exoplanet τ Boötis b via HRS with VLT/CRIRES. Despite the two signals have approximately the same order of magnitude, the host M star would be at least 6 magnitudes fainter, requiring an ELT-like telescope. Adapted from [89]

5.1.1 Detecting oxygen in planets around M dwarfs

Figure 72, adapted from [89], shows the high-resolution transmission spectrum of the Oxygen-A band for an Earth orbiting an M5V star, in comparison to the CO signal that was detected from the hot Jupiter τ Boötis b [16]. Due to the considerations made in Section 5.1 applied to the concepts of Section 1.5.1 about how transmission signals scale, the average line depth of the O_2 lines is about 1/3 of the CO signal from τ Boo b. However, the challenge comes when taking into account that even the closest M5 stars are much fainter than τ Boo. In the *I* band, which is a good match to the oxygen A-band, even the closest M dwarfs will be at least 6 magnitudes fainter, which means that in order to reach the appropriate S/N Extremely Large Telescopes are needed and tens of transits co-added.

Going a bit deeper than a straightforward scaling of the signal, it has been shown that instruments with high efficiency and observations optimised to have the maximum radial-velocity separation between the telluric oxygen and the exoplanet spectrum are both key to a successful detection [48, 76, 82]. Therefore, while feasible in theory, it might take a few tens of transit (or equivalently a few years of carefully planned observations) with one of the next-gen giant telescopes (ELT, GMT, or TMT) to reach enough S/N to claim a detection. Nevertheless, it would be the first detection of a potential biomarker outside of the solar system, which is quite an exciting prospect.

5.1.2 Challenges in M-dwarf observations

M dwarfs pose significant challenges to the habitability of terrestrial exoplanets related to their environments, such as activity, flaring, luminosity changes in the pre-main-sequence phase, and the overall spectral energy distribution peaking in the near-infrared [83]. Furthermore, due to the proximity to the parent star, any planets in the HZ would be likely tidally locked, with unknown consequences on their climate.

Here we will not focus on the physical processes challenging habitability around M dwarfs, but rather list some of the main observational challenges. Firstly, the faintness of these stars in the optical band make any search for reflected light, which would be favourable in terms of planet/star contrast (see Section 1.5.2), hard in practice. In fact simple back of

the envelope calculations for Proxima Cen b (which once again is the closest M dwarf to the solar system) show that we would need 100 hours of ELT/GMT time to detect such a planet even considering that M dwarf spectra are very rich in spectral lines and a gain of 65-80 in S/N can be easily obtained from cross correlating with their spectra.

In addition, the same complexity of their spectra is also a limiting factor, in that M-dwarf show molecular lines - such as TiO or H₂O, which are not accurately known for high-resolution models, or can substantially correlate with the signals of exoplanets.

Thirdly, while stellar variability is not a big limit to HRS analysis if it only affects the continuum, it is well known that flares — frequent and sudden increases in luminosity — affect particularly some spectral lines (for instance the H- α line), potentially complicating or even invalidating the application of the analysis shown in Section 2.3. Some research groups are actively tackling the effects of flares at the time of writing, with promising preliminary results (J. Birkby, private communication).

Lastly, M dwarf have generally strong magnetic fields, and this has an influence on the frequency of their spots and the shape of their spectra. Correcting these late-type spectra through coupled 3D MHD simulations of stellar photospheres similar to current hydro-dynamic simulations shown in Section 3.4 is still beyond our current capabilities, and will arguably require significant work in the future.

5.2 Combining spectral and spatial resolution

Direct imaging of exoplanets is the closest technique that we can imagine to “taking a picture” of an exoplanet system. It consists in separating *spatially* the photons from the star and those from the planet. This task is far from trivial, as due to the enormous distances involved a planet-star system is separated by a tiny angle on the sky as seen from the Earth, thus requiring big telescopes and adaptive optics systems in order to correct for the distortions caused by the Earth’s atmosphere. Furthermore, planets are also orders of magnitude fainter than their parent stars, requiring sophisticated techniques to mask out the enormous glare of the star and reveal the planet.

Current technology can detect exoplanets on wide orbits that are up to a million times fainter than the star, and at a fraction of an arc-second of angular distance. However, as this is still tens of AUs at the distances of the nearest bright stars, none of the evolved exoplanets would be detectable, because even the biggest giants would be either too cold or too far to show any reasonable signals given the current technology (compare e.g. with the estimates we obtained for thermal emission and reflected light in Sections 1.5.3 and 1.5.2).

The workaround is to target young exoplanet systems. These contain giant planets that are still contracting from their formation, and thus they have significant internal energy left - which means a much higher temperature than at equilibrium with the stellar radiation, as high in fact as that of an evolved hot Jupiter. For some of the young directly imaged systems, e.g. β Pictoris or HR 8799, we have become so good at monitoring these planets that we can see them moving along the orbit throughout the years.

As promising as direct imaging is, our current sensitivity and technology implies that the scales and signals of planets in our solar system would be completely out of range. While technology is being developed that would lead to the required contrasts of 10 to 100 billion between the star and the planet, in recent years a hybrid methodology has been proposed combining *spatial* and *spectral* resolution [50, 54, 87, 90, 97]. In the context of this Section this technique will be abbreviated as HRS+HCI, from High-Resolution Spectroscopy and High-Contrast Imaging.

To understand the principles of this novel observing strategy, it is best to consider again the signal-to-noise formula (Equation 47). In Section 2.2 this has been manipulated to show

that cross correlation can amplify the signal by a factor of $\sqrt{n_{\text{lines}}}$ (Equation 49). It has been also pointed out that for bright targets, being star and planet spatially unresolved, the photons from the star $N_{\star}$ dominate the noise budget, i.e. the denominator. It follows that if we can suppress the photons from the star *at the position of the planet* by a factor of K via HCI, i.e. by applying direct imaging techniques, this would reduce the corresponding noise budget by a factor $\sqrt{K}$, thus increasing the overall S/N by the same factor. The S/N formula for an HRS+HCI observation would therefore be

$$\left(\frac{S}{N}\right)_{\text{CCF+HCI}} = \sqrt{n_{\text{lines}}} \frac{nN_{\text{p}}}{\sqrt{N_{\star}/K}}, \quad (58)$$

at least until the photons recorded from the star dominate over the background and the electronics.

[90] put in practice the above concepts, and to do so they used again the CRILES spectrograph at the Very Large Telescope. Although this is not a dedicated direct-imaging machine, it has an adaptive optics system attached that somewhat brings the spatial resolution of the telescope closer to its nominal value for an 8.2-metres telescopes. The planet targeted was β Pictoris b, at the time of the observations 0.7'' away from its host star. The star and the planet were aligned along the spatial direction of the instrumental slit and a contrast of $K = 7 - 20$ was achieved at this angular distance. The data was then cross correlated with models for the planet atmosphere, which would be a poor match to the stellar photosphere due to the early spectral type of β Pic (an A-type star). The resulting two-dimensional cross correlation (one spatial direction along the slit, and one radial-velocity direction perpendicular to it) peaked at the expected position of the exoplanet and significantly above the noise level, which allowed the authors to derive information about the planet's atmosphere, including its rotational rate. Similar success [81] with the same instrument was obtained for another directly-imaged planet (GQ Lupi b), until CRILES was finally decommissioned in Summer 2014.

A more suited instrument for HRS+HCI would be an *Integral Field Unit* (IFU), that is an instrument that can take a picture of a small region of the sky, and also a spectrum for each pixel of the image. At the moment the only IFUs available are limited to a spectral resolving power of at most 5,000. This is somewhat sufficient to resolve the strongest molecular features of some absorption bands, and thus it was used with cross correlation techniques to achieve detections for β Pictoris b again [36] and HR 8799 b [44].

The good news is that one of the earlier instruments mounted at the Extremely Large Telescope (METIS) will have IFU capabilities at a resolving power of 100,000, thus unlocking the full potential of the method in combination with the large aperture of the telescope. Integration of HRS+HCI at the next-generation of ground-based facilities is the subject of many active lines of research, and most of the information is still circulating at the level of preliminary study, which makes it unsuitable to be reported in this Chapter. However, a reasonable expectation is that an ELT with performing IFUs in the near-infrared should yield a detection of thermal emission from an Earth orbiting α Centauri B in a few tens of hours of observations [87], and the same observing time would yield in the optical a detection of reflected light from the known Proxima Centauri b.

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Stellar Winds and Planetary Atmospheres

by C. Johnstone



Artist's impression of the effect of stellar radiation and wind on planetary atmosphere. Copyright: NASA's Goddard Space Flight Center. Licensed under the Creative Commons Attribution-Share Alike 4.0 International. Source: https://commons.wikimedia.org/wiki/File:HD_189733b%27s_atmosphere.jpg

Stellar Winds and Planetary Atmospheres

Colin Johnstone^{1,2,*}

¹Natural History Museum, Burgring 7, A-1010 Vienna, Austria

²University of Vienna, Department of Astrophysics, Türkenschanzstrasse 17, 1180 Vienna, Austria

Abstract. Interactions between the winds of stars and the magnetospheres and atmospheres of planets involve many processes, including the acceleration of particles, heating of upper atmospheres, and a diverse range of atmospheric loss processes. Winds remove angular momentum from their host stars causing rotational spin-down and a decay in magnetic activity, which protects atmospheres from erosion. While wind interactions are strongly influenced by the X-ray and ultraviolet activity of the star and the chemical composition of the atmosphere, the role of planetary magnetic fields is unclear. In this chapter, I review our knowledge of the properties and evolution of stellar activity and winds and discuss the influences of these processes on the long term evolution of planetary atmospheres. I do not consider the large number of important processes taking place at the surfaces of planets that cause exchanges between the atmosphere and the planet's interior.

1 Introduction

The discoveries in the last decades of thousands of planets outside our solar system has led to questions of planetary habitability becoming a central theme in modern astrophysics. Central to these questions is the question of how planetary atmospheres evolve, and what conditions are necessary to form an atmosphere that can sustain life (for a review, see [71]). The formation and subsequent evolution of a planet's atmosphere are determined by exchanges of volatiles between the planet's interior and atmosphere and by losses of gas from the system to space by a large number of processes. These loss processes take place mostly in response to magnetic activity related output of the planet's host star, and especially its emission of X-ray and ultraviolet radiation and its wind. In this chapter, I discuss the properties and evolution of stellar winds and the various ways in which they influence planetary atmospheres. For a more comprehensive review on the topic, see the recent book by [122].

Atmospheres form on planets in several ways and different processes lead to atmospheres composed of different mixtures of chemical species. Planets that form to sufficient mass (typically above $0.1 M_{\oplus}$) in the first few million years when a significant gas disk is still present can accumulate hydrogen and helium dominated primordial atmospheres and the amount of gas accumulated depends primarily on the mass of the planet at the end of the disk phase ([181]). [115] showed that since higher mass planets accumulate more atmosphere and lose them slower after the disk is gone, planets with masses above approximately an Earth mass keeps these atmospheres for their entire lives and a similar result was found for planets orbiting M dwarfs by [149]. Observations of low-mass exoplanets with large radii (e.g. [123];

*e-mail: colin.johnstone@univie.ac.at

[167]) suggest that many terrestrial planets have such envelopes, and especially higher mass planets (e.g. [167]). Evidence from the isotopic abundances of noble gases in the lower mantle suggest that the Earth did in fact possess a primordial atmosphere ([215]; [205]), though disagreement exists about how to interpret these results ([158]). The isotopic compositions of hafnium and tungsten in the mantle suggest that the Earth likely grew to half its mass in approximately 10 Myr and its full mass within approximately 50 Myr ([108]). Approximately half of all gas disks are thought to be gone by ~ 2 Myr ([127]), though there is evidence that a significant minority of stars retain their disks for much longer than 10 Myr ([159]). It is plausible that the Earth gained a small primordial atmosphere, though since this atmosphere is no longer present, it likely did not form to its full mass during the gas disk phase.

Volatile species are distributed throughout several reservoirs in a planet, including the mantle, the crust, the ocean, and the atmosphere. Impacts during a planet's growth phase lead to volatile species being released into the atmosphere from both the impactor and the planet, and what is not released from the impactor can be stored within the planet outer layers for release at later times ([52]). Energy released from these impacts likely cause the planets to be so hot that a liquid magma ocean is present during and immediately after the main growth period. As the planet solidifies and forms a mantle and crust, large amounts of volatiles can be degassed into the atmosphere ([51]; [168]) and if it is not rapidly lost to space, this atmosphere can lengthen the magma ocean phase by slowing the radiative cooling of the surface ([144]). In later phases, tectonic related activity, such as mantle outgassing from volcanoes, is important for forming and replenishing atmospheres (e.g. [146]), which can even be directly influenced by stellar magnetic fields ([103]; [104]). Surface processes can also transfer gas into the interior and this can be more important for removing atmospheres than escape to space.

The fact that losses to space can play a major role in an atmosphere's evolution can be seen in our own solar system. Venus for example has likely undergone massive loss of water due to the dissociation and loss of water molecules in the upper atmosphere ([27]). Since hydrogen atoms are lost more efficiently than the heavier deuterium atoms, this massive water loss can be seen in Venus' atmosphere by the very large D/H ratio compared to other solar system bodies ([129]). While Mars currently has a very thin atmosphere, there is strong evidence that it previously had liquid water on its surface ([107]) meaning that it must have had a much thicker atmosphere that has since been lost. This is consistent also with the atmosphere's isotopic composition which contains enhanced abundances of heavier isotopes of certain elements relative to their lighter counterparts ([83]). Outside our solar system, atmospheric escape has been directly observed from exoplanets ([50]) and this escape likely has observable influences on the distributions of exoplanet radii as measured by spacecraft such as *Kepler*. For example, the 'evaporation valley' has been predicted to result from such losses and has significant observational support ([150], [152]; [61]).

A star's magnetic field is one of the most important factors for the structure of its outer atmosphere and causes the surface of the star to be inhomogeneous and time variable, which can complicate the detection and characterisation of exoplanets ([34]). Field lines anchored in the photosphere where the plasma beta (the ratio of the thermal pressure to the magnetic pressure) is much greater than unity are moved around by convective motions of the plasma. Although it is not yet clear what the dominant processes are, energy from these convective motions is transferred by the magnetic field to higher altitudes where it heats the gas to several tens of thousands of K in the chromosphere and to MK temperatures in the corona. The corona is the large outer atmosphere of the star and has a plasma beta much less than unity, meaning that the dynamics of the plasma is controlled by the magnetic field. This results in the emission of X-ray and extreme ultraviolet (EUV) radiation from the corona, and EUV and far ultraviolet (FUV) radiation from the chromosphere. The term 'XUV' is commonly used to

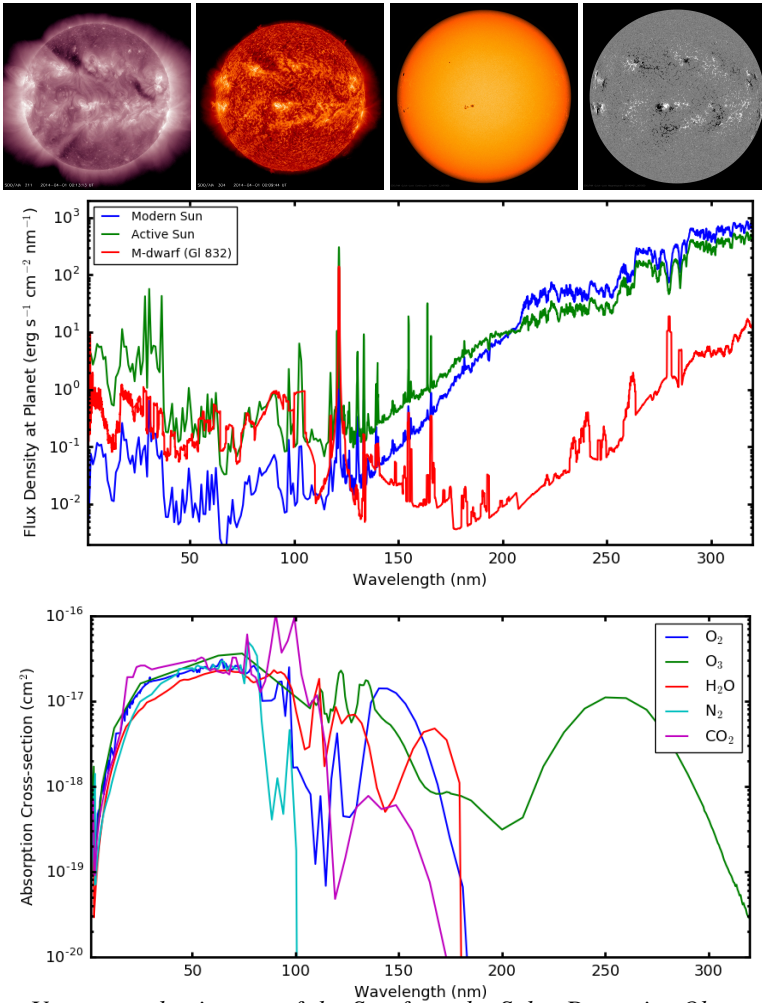


Figure 73: *Upper panels: images of the Sun from the Solar Dynamics Observatory courtesy of NASA/SDO and the AIA, EVE, and HMI science teams. From left to right, the panels show the corona at 2×10^6 K, the chromosphere and transition region at 5×10^4 K, the photosphere, and the surface magnetic field structure (with black and white showing opposite polarity magnetic field). Middle panel: typical XUV spectra of the modern Sun during activity maximum (blue line) and a more active young Sun (green line) from ([29]), and the low-mass M-dwarf Gliese 832 as modelled by [59]. The solar spectra are scaled to 1 AU and the Gliese-832 spectrum is scaled to 0.162 AU, which is the orbit of a known habitable zone planet. The lower flux of the Gliese 832 spectrum longward of ~ 150 nm is due to the much cooler photosphere of this star, meaning its photospheric spectrum is shifted to longer wavelengths. Note that the resolution of the Gliese-832 spectrum has been reduced to be similar to the other two spectra. Lower panel: XUV absorption cross-sections for several important species in planetary atmospheres from the PHIDRATES database ([81]).*

refer to a star's X-ray and ultraviolet spectrum, though definitions of this term vary and often includes only X-ray and EUV radiation; in this review, I do not attach an exact definition to the term unless specified. Another effect of these high temperatures is the acceleration of the plasma away from the star in the form of a transonic (starting subsonic and accelerating to

supersonic speeds) wind. The outer layers of the Sun's atmosphere and the resulting X-ray and ultraviolet spectrum of the modern Sun, a younger more active Sun, and a low mass M dwarf, are shown in Fig. 73. This high energy radiation is important since most common atmospheric species absorb radiation at these wavelengths very efficiently, meaning that the radiation is absorbed high in the atmospheres of planets where the gas densities are low. This heats and expands the upper atmosphere, making it more susceptible to interactions with the star's wind.

2 Stellar rotation and activity evolution

It is important to consider the long term evolution of stellar rotation since we should expect that this will drive a corresponding evolution in stellar winds and their interactions with atmospheres. Possibly the most important way that winds influence the atmospheres of planets is by removing angular momentum from their host stars, causing a decay in both their rotation rates and their emission of X-ray and extreme ultraviolet radiation. In the last two decades, our observational knowledge of rotational evolution has improved dramatically, especially due to rotation distributions in young stellar clusters being determined by ground-based photometric monitoring campaigns (e.g.[76]) and more recently by the *K2* mission (e.g. [161]). This has driven the development of more detailed physical models and a comprehensive though still incomplete understanding of how the rotation rates of stars evolve from their initial birth to the end of the main-sequence.

2.1 The observational picture

Stars in young clusters with ages of less than 5 Myr have a range of rotation rates distributed between approximately 1 and 50 times the solar rotation rate ([4]). This distribution appears to be initially mass independent and during the first few Myr, it remains approximately constant in time ([162]), though there is evidence that stars with masses below approximately $0.5 M_{\odot}$ might spin up during this time ([78]). At an age of 12 Myr, the rotation distribution in young cluster h Per studied by [139] shows two changes relative to younger clusters. Firstly, the distribution has been shifted to more rapid rotation, especially for the rapidly rotators. Secondly, the distribution has separated into a rapidly rotating track and a slowly rotating track, with few stars in between. This distribution can be seen in Fig. 74 and the evolution of this distribution on the main-sequence can be seen by comparison with the Pleiades (~ 130 Myr) and Hyades (~ 650 Myr) clusters. For Pleiades, rotation rates from [76] and [161] give us a detailed description of the rotation at this important age. At this age, most stars with masses above approximately $0.5 M_{\odot}$ are part of the slowly rotating track, but a tail of rapid rotators is also present, while most lower mass stars remain rapidly rotating. By 650 Myr, this distribution has shifted to slower rotation and more stars have converged onto the slowly rotating track, though this convergence is slower for lower mass stars. This spin-down and convergence continues throughout their main-sequence evolution, though recent measurements of K dwarfs seem to show a temporary halt in their spin-down between ages of approximately 500-1000 Myr ([42]). Although at young ages they remain rapidly rotating longer, at later ages lower mass stars are more slowly rotating than higher mass stars ([143]). [82] provided evidence that some fully convective M dwarfs do not converge onto the slowly rotating track and remain rapidly rotating for timescales longer than 10 Gyr.

While observations constrain the first Gyr of rotational evolution, how stars spin down after this time is less well constrained. Especially due to *Kepler* measurements, we know the rotation rates of a very large number of field stars ([134]; [169]), which are informative but of

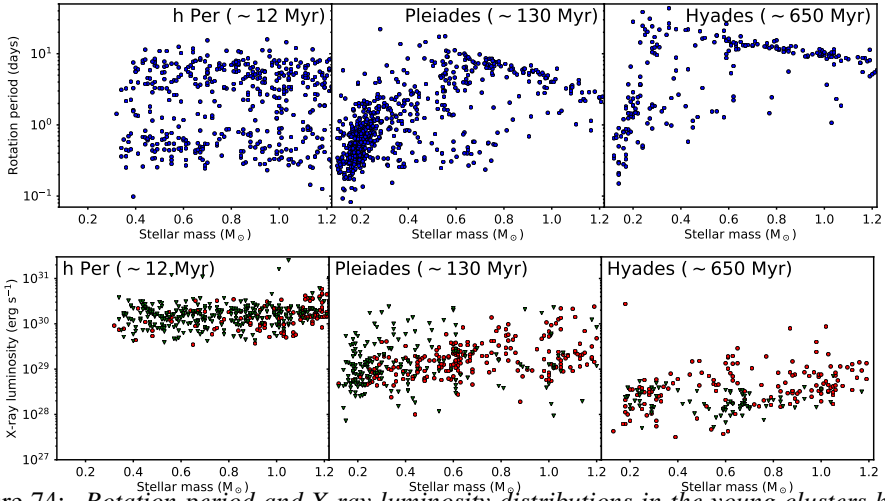


Figure 74: *Rotation period and X-ray luminosity distributions in the young clusters h Per, Pleiades, and Hyades. Rotation periods are from [139] for h Per, [76] and [161] for Pleiades, and [44] for Hyades. The X-ray luminosities are from Argiroffi et al. (2016) for h Per, and Núñez et al. (2016) for both Pleiades and Hyades.*

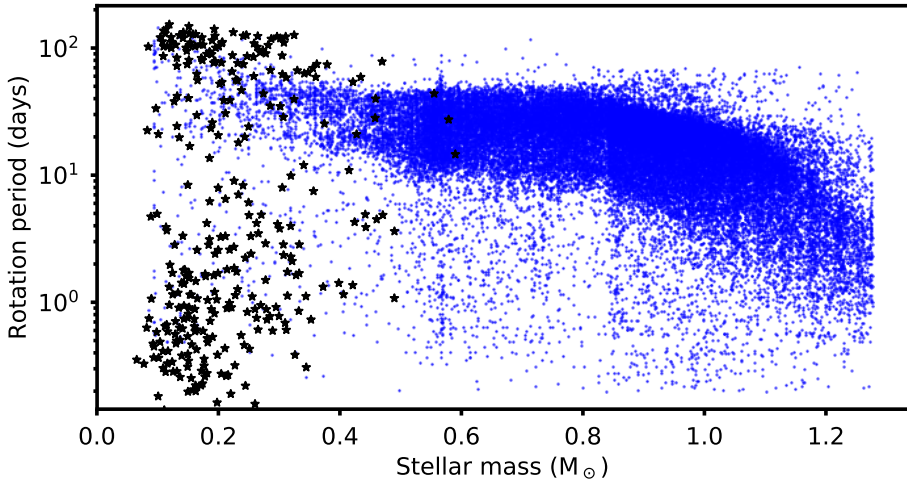


Figure 75: *Rotation distribution for approximately 40,000 field stars from [134] and [169] measured by Kepler (blue points) and for low mass field stars from [142] by the MEarth Project (black stars).*

more limited use since the ages of these stars are difficult to determine ([177]). In Fig. 75, I show the rotation-mass distribution for field stars measured by the *Kepler* spacecraft and for lower mass stars by [142]. At later ages, higher mass stars tend to be rotating more rapidly, which could be surprising given that in the first billion years, lower mass stars remain rapidly rotating longer. It is also shown from the MEarth Project sample of lower mass stars that even for field stars, most of which likely have ages of a few Gyr or more, there are still a large number of very rapid rotators.

The rotational evolution of solar mass stars was described in by [176] by $P_{\text{rot}} \propto t^{0.5}$, where P_{rot} is the rotation period and t is the age. While this was originally based only on comparing

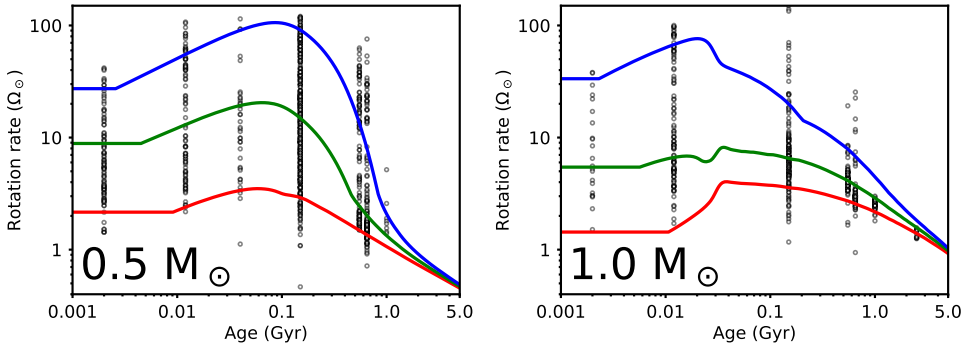


Figure 76: *Evolutionary tracks for surface rotation of stars with masses of 0.5 and 1.0 $M_{\odot}$ with different initial rotation rates. In both panels, the red, green, and blue lines show stars at the 5th, 50th, and 95th percentiles of the observed rotation distribution at 150 Myr and the black circles show observed rotation rates from several young clusters. The rotation tracks are calculated using the model developed by [95] and extended by Johnstone et al. (submitted).*

average equatorial rotation velocities from Hyades and Pleiades with the rotation of the Sun, this relation has been shown to be approximately accurate for solar mass stars at later ages, though this conclusion is dependent on the assumption that the Sun’s modern rotation is a good proxy for the rotation rates of solar mass stars with similar ages. The mass dependence of rotation at later ages is more difficult to constrain and it is typical to assume that P_{rot} depends both on mass and age and to constrain the mass dependence using the shape of the slow rotating track in young clusters (e.g. [12]; [128]). The reliability of this assumption is difficult to test given the lack of reliable ages for older stars with masses below that of the Sun. Based on this assumption, [128] derived

$$P_{\text{rot}} = a [(B - V)_0 - c]^b t^n \quad (59)$$

where $a = 0.407$, $b = 0.325$, $c = 0.495$, $n = 0.566$, t is in Myr, and P_{rot} is in days. Here, $(B - V)_0$ represents stellar mass and solar mass main-sequence stars have values of approximately 0.66, whereas 0.5 and 0.1 $M_{\odot}$ stars have values of approximately 1.5 and 2.0 respectively ([157]).

2.2 The epochs of rotational evolution

A star’s long term rotational evolution depends primarily on its mass and its early rotation rate ([63]). The latter is important because stars that are born as rapid rotators remain rapidly rotating much longer than those born as slow rotators. Other important factors include the star’s metallicity ([7]) and possibly the lifetime of its circumstellar disks during its classical T Tauri phase ([79]). A star’s rotational evolution from 1 Myr to the end of the main-sequence can be broken down into several important epochs: these are the phases of disk-locking, pre-main-sequence spin-up, and main-sequence spin-down. The rotational evolution of 0.5 and 1 $M_{\odot}$ stars is shown in Fig. 76.

The observed lack of rotational spin up for classical T Tauri stars in the first few million years is surprising since they are contracting on the pre-main-sequence and are accreting high angular momentum material from their disks. This means that angular momentum must be removed from the stars and potentially also from the accreting material during this time,

though which mechanisms are responsible are poorly understood. A possibility is that very strong winds originating from the surfaces of the stars exert a strong enough spin-down torque to keep the stars in rotational equilibrium ([131]) and this strong wind has been suggested to be powered by accretion onto the surface from the gas disk ([36]; [38]). This phase is commonly referred to as ‘disk-locking’ given the assumption that interactions with the disk are responsible (for a review, see [21]) and it is generally assumed that this phase ends around the time that these disks disperse, which is typically within the first 10 Myr ([127]), after which stars spin up due to pre-main-sequence contraction. This spin-up is reduced by the removal of angular momentum by stellar winds and stops approximately at the zero-age main-sequence (ZAMS).

After the ZAMS, the main process is wind driven spin-down. Several factors are important for how rapidly winds remove angular momentum from their host stars and these are mostly related to the star’s mass, radius, rotation rate, and the properties of the wind and global magnetic field ([97]; [130]). Characterising the dependence of angular momentum loss on these parameters is difficult and recent magnetohydrodynamic models have been used to explore this question (e.g. [133]). Important is not only the strength of the magnetic field, but also its geometry ([64]; [57]) and thermal structure ([31]; [154]).

As I discuss in the next section, stellar activity has a saturated (activity independent of rotation) regime and an unsaturated (activity dependent of rotation) regime. Observationally, it appears that main-sequence stars in the saturated regime show exponential spin-down which implies that $\dot{\Omega} \propto \Omega$, where $\dot{\Omega} = d\Omega/dt$, whereas the Skumanich spin-down of unsaturated stars implies $\dot{\Omega} \propto \Omega^3$. This difference can be easily understood if mass loss rates and magnetic field strengths are independent of rotation in the saturated regime and depend strongly on rotation in the unsaturated regime. In both regimes, the positive dependence of $\dot{\Omega}$ on Ω means that the wide distribution of rotation rates seen at the start of the main-sequence converges until all stars with the same mass follow a single spin-down track. Activity saturation also explains why rapidly rotating lower mass stars spin down slower than rapidly rotating higher mass stars, despite the fact that at later ages, lower mass stars spin down more rapidly. Since lower mass stars saturate at lower rotation rates, in the saturated regime they feel a reduced spin-down wind torque. In the unsaturated regime, wind torques are approximately mass independent ([63]; [90]), meaning that lower mass stars spin down more rapidly due to their smaller moments of inertia.

2.3 Activity and rotation

Empirically, it is known that the activity of a star depends primarily on its mass, age, and rotation rate. This dependence on rotation can be seen in direct measurements of photospheric magnetic fields ([196]), in chromospheric activity indications such as Ca II H&K and H α lines ([19]), and in coronal X-ray emission ([160]). For most activity indicators, the rotation relation has two regimes. At slow rotation a strong rotation–activity relation is seen with more rapidly rotating stars being more active and this regime is called the unsaturated regime. At fast rotation, the rotation–activity relation is very flat with almost no difference between stars with different rotation rates and this regime is called the saturated regime. On the pre-main-sequence, the rotation rate of the saturation threshold is initially very low and increases rapidly. This means that at very young ages (younger than maybe 10 to 20 Myr), rotation is not a relevant factor for determining a star’s activity. For example, the X-ray luminosities of stars in very young clusters depend on mass and age ([70]; [69]). On the main-sequence, the saturation threshold is at approximately $15\Omega_{\odot}$ for solar mass stars and is at lower rotation rates for lower mass stars. This means that low mass stars must reach slower rotation

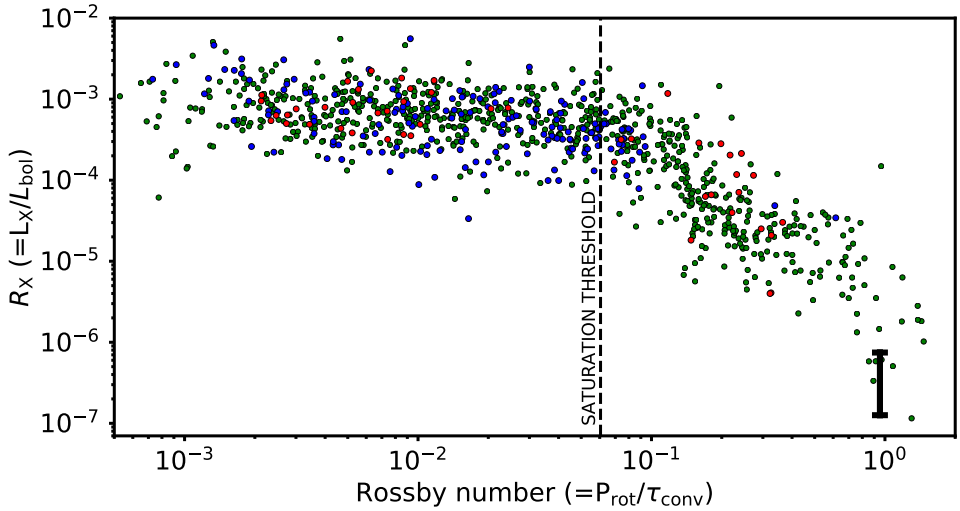


Figure 77: The correlation between stellar X-ray emission and rotation characterised as a Ro – R_X relation. Here, green circles show main-sequence stars from [212], blue circles show stars in the young 12 Myr old cluster h Per from [10], and red circles show fully convective main-sequence M dwarfs from [86] and [213]. I do not include upper limits from these studies and for main-sequence M dwarfs these upper limits could indicate that the slope in the unsaturated regime is steeper than for higher mass stars. The black line on the lower right shows the range of values for the modern Sun (specifically the 10th to the 90th percentiles) calculated from the Flare Irradiance Spectrum Model ([24]). All values of Ro have been calculated using the convective turnover times from [179].

rates before spin-down related activity decay starts, which contributed to lower mass stars remaining highly active longer.

The dependence of X-ray emission on stellar rotation, mass, and age are well represented as a dependence between the X-ray luminosity normalised to the bolometric luminosity, $R_X = L_X / L_{bol}$, and the Rossby number, Ro , which is a dimensionless parameter defined as $Ro = P_{rot} / \tau_c$, where P_{rot} is the rotation period and τ_c is the convective turnover time. Note that several studies have discussed the possibility that using rotation period instead of Rossby number is better for describing the activity–rotation relation ([163]; [126]). The Ro – R_X relation is shown in Fig. 77 and can be described as a broken power-law given by

$$R_X = \begin{cases} c_1 Ro^{\beta_1}, & \text{if } Ro \geq Ro_{sat}, \\ c_2 Ro^{\beta_2}, & \text{if } Ro \leq Ro_{sat}, \end{cases} \quad (60)$$

where Ro_{sat} is the saturation Rossby number and c_1 , c_2 , β_1 , and β_2 are constants that can be determined empirically. If we assume that R_X is constant in the saturated regime, as is common in the literature, then $\beta_1 = 0$ and $c_1 = R_{X,sat}$. Making this assumption, [212] found $\beta_2 = -2.18$, $Ro_{sat} = 0.13$, and $R_{X,sat} = 10^{-3.13}$, though depending on the method used for fitting in the unsaturated regime, values for β_2 between ~ 2 and 3 are often found ([163]). A weak dependence between Ro and R_X can be seen in the saturated regime too and there is evidence from extremely rapidly rotators for ‘supersaturation’, which is a decrease in X-ray emission with increasing rotation at the high rotation part of the saturated regime ([84]; [10]).

It is not yet fully clear how well pre-main-sequence and fully convective main-sequence stars follow the same Ro – R_X relation as higher mass main-sequence stars. In Fig. 77, I show the Ro – R_X relation for both pre-main-sequence and main-sequence stars, including

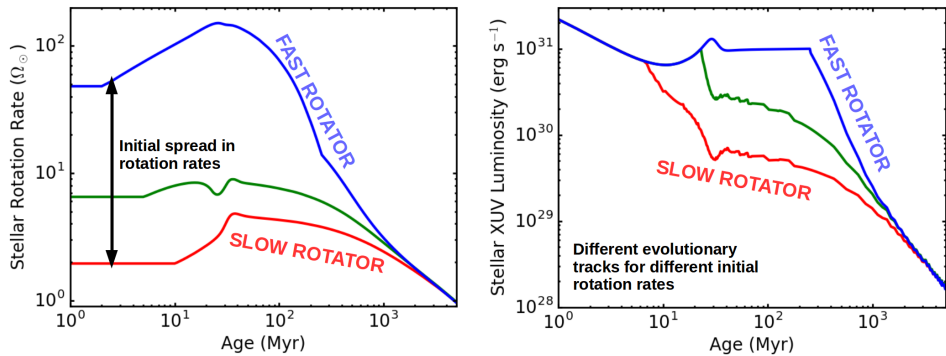


Figure 78: *Evolutionary tracks for rotation (left) and XUV emission (right) from [189] for solar mass stars with different initial rotation rates. Here, XUV is calculated considering just X-ray and EUV wavelengths.*

for main-sequence stars also fully convective M dwarfs in the unsaturated regime. For pre-main-sequence stars, the question is difficult to answer because in most very young stellar clusters, most stars are saturated regardless of their rotation rate due to the very large convective turnover times. [10] studied the X-ray–rotation relation in the 12 Myr old cluster h Per which is useful because the age of this cluster means that many of the slowest rotators could be unsaturated. Their results suggest a much shallower relation between R_o and R_X (meaning a smaller β_2) at this age in the unsaturated regime though this depends on the convective turnover times used and Fig. 77 suggests that the R_o – R_X relation in h Per is consistent with the main-sequence relation. For fully convective main-sequence M dwarfs, the difficulty is similarly that the long convective turnover times mean that all but the very slow rotators are in the saturated regime making it hard to characterise the unsaturated regime. This was studied for the young open cluster NGC 2547 by [86] who found that the unsaturated regimes for fully convective M dwarfs and higher mass stars are similar as can be seen in Fig. 77, and several studies have backed up their conclusions ([211]; [213]). There is evidence that the slope for fully convective stars is steeper ([126]) which might not be visible in Fig. 77 since I do not include upper limits.

2.4 Activity evolution

Given the close link between rotation and activity, we should expect that the activity levels of stars decay with age due to rotational spin-down. This was observed by [176] who found that Ca II emission decays simultaneously with rotation over billions of years. For X-ray emission from solar mass stars, this decay was characterised by [72] who found $L_X = 2.1 \times 10^{28} t^{-1.5} \text{ erg s}^{-1}$, where t is the age in Gyr. This decay in high energy emission is seen for all magnetic activity related emission, including extreme ultraviolet, though at longer wavelengths the decay law is less steep ([165]; [29]). This is in part due to the fact that more active stars have hotter emitting plasma ([171]; [89]), meaning that the spectrum is more shifted to shorter wavelengths. If we consider again the link between rotation and high-energy emission and the various possible rotational evolution tracks that stars can follow (Fig. 76), we can see that there is a problem with describing activity decay as a single unique power-law. Since stars born as fast rotators remain rapidly rotating for much longer than those born as slow rotators, we should expect that the initial rotation rate of a star plays an

important role in determining the evolution of its emission. This was shown in [189] for solar mass stars, and three different possible evolutionary tracks for X-ray and EUV emission from this study are demonstrated in Fig. 78. The slow, medium, and fast rotator tracks correspond to the 10th, 50th, and 90th percentiles of the rotation distribution and the evolution of X-ray emission (0.1–10 nm) for these three tracks can be approximated as

$$L_X = \begin{cases} 2.0 \times 10^{31} t^{-1.12} \\ 2.6 \times 10^{32} t^{-1.42} \\ 2.3 \times 10^{36} t^{-2.50} \end{cases}$$

where L_X is in erg s⁻¹ and t is in Myr. The evolution of the EUV emission (10–91 nm) for these three tracks can be approximated as

$$L_{\text{EUV}} = \begin{cases} 7.4 \times 10^{31} t^{-0.96} \\ 4.8 \times 10^{32} t^{-1.22} \\ 1.2 \times 10^{36} t^{-2.15} \end{cases}$$

where the units of the quantities are the same as the previous equations.

3 Properties and Evolution of Stellar Winds

In the previous section, I concentrate on the evolution of stellar rotation and its connection to magnetic activity without discussing in any detail winds, which is a primary topic of this review. In this section, I discuss how stellar wind bulk properties – primarily mass flux and outflow speed – evolve with time and how this depends on the parameters of the star, such as mass and initial rotation rate. Unfortunately little is known about this for low-mass stars other than the Sun and what is known is not known with any certainty.

3.1 Observational constraints on winds

While observations of the long term evolution of stellar rotation make it clear that winds are ubiquitous among low-mass stars, other than for the case of the solar wind, we do not possess clear observational constraints on their properties. Much work has been dedicated to detecting and characterising winds but since winds consist of very tenuous gas, this is a very difficult task. Most methods that have been developed attempt to indirectly determine wind properties by measuring the effects of these winds on the local interstellar environment around the star, planetary or stellar companions, or the star itself.

3.1.1 Rotational evolution

The rotational evolution of stars is the most clear indication that stellar winds are present on low-mass stars and possibly has the most potential for allowing us to learn about wind properties. For example, the observed differences between the spin-down rates of saturated and unsaturated stars discussed in Section 2.2 is a good indication that the mass loss rates of winds saturate at rapid rotation with other activity related phenomena such as X-ray emission. The rate at which a star spins down on the main-sequence depends on several factors, including its mass and radius, the properties of its magnetic field, and the properties of its wind and it is therefore possible to determine wind properties from the observed rotational evolution of stars as explored by several studies ([62]; [131]). The model of [204] for angular momentum loss from the winds of magnetised rotating stars gives the angular momentum loss from a

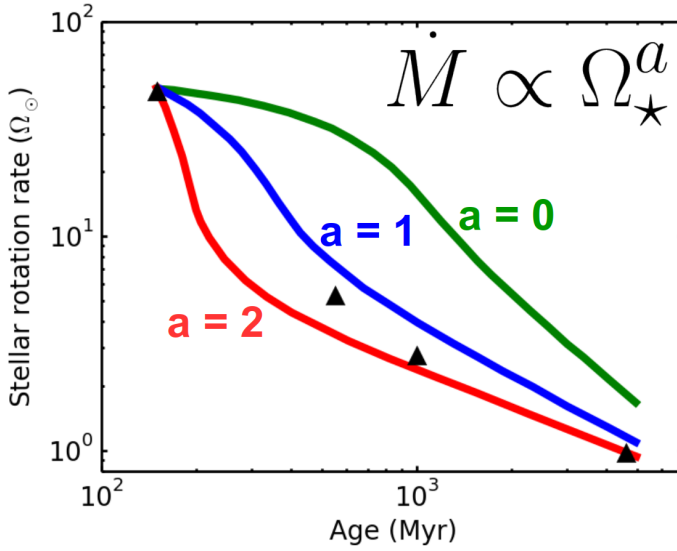


Figure 79: Figure demonstrating how the observed evolution of stellar rotation can be used to constrain the properties of stellar winds based on the rotational evolution model of [90]. The black triangles from show observational constraints on the evolution of stars at the 90th percentile of rotation distributions for solar mass stars. The rotation tracks start at the first constraint and show cases for different values of a in the assumed equation $\dot{M}_w \propto \Omega_\star^a$. The best fit value of a was found to be 1.33 for this particular rotational evolution model.

stellar wind as $\tau_w \propto \dot{M}_w \Omega_\star R_A^2$, where $\dot{M}_w$ is the wind mass loss rate, $\Omega_\star$ is the stars rotation rate, and R_A is the Alfvén radius. Past ages of approximately 1 Gyr, the spin-down of solar mass stars can be approximated as $\Omega_\star \propto t^{-0.5}$ ([176]), where t is the age, implying that $\dot{\Omega}_\star \propto \Omega_\star^3$. If we assume approximately constant values for the stellar moment of inertia and the Alfvén radius, this implies that $\dot{M}_w \propto \Omega_\star^2 \propto t^{-1}$. While this reasoning is overly simplistic, it demonstrates the potential to derive wind properties from observed rotational evolution.

Ideally, we would determine wind properties using a full detailed rotational evolution model fit to the abundant observational constraints on rotation distributions in young stellar clusters. Using such a model, [132] found that $\dot{M}_w \propto M_\star^{1.3} Ro^{-2}$, where Ro is the Rossby number. Similarly, considering only stars after the first 100 Myr, [90] found $\dot{M}_w \propto R_\star^2 \Omega_\star^{1.33} M_\star^{-3.36}$. A plausible assumption is that the surface mass flux has power-law dependencies on Rossby number and mass, which gives

$$\dot{M}_w \propto R_\star^2 Ro^a M_\star^b, \quad (61)$$

where a and b are parameters to be determined. Note that the above is only when Ro is greater than the saturation Rossby number, Ro_{sat} , and when this is not the case, Ro_{sat} should be used in place of Ro . Johnstone et al. (submitted) used measurements of stellar rotation and activity to constrain these parameters and found $a = -1.76$, $b = 0.649$, and $Ro_{\text{sat}} = 0.0605$. This equation can be used to find the mass loss rate of any star in the mass range 0.1 to 1.2 $M_\odot$ by using the solar mass loss rate of $\dot{M}_\odot = 1.4 \times 10^{-14} M_\odot \text{ yr}^{-1}$ and the convective turnover times derived by [179], though there is much uncertainty for the determination of these parameters. This suggests that $\dot{M}_w \propto t^{-0.88}$ for the solar wind after the first Gyr. The estimates for $\dot{M}_w$ given here are possibly the best observationally driven constraints on wind

mass loss rates available, but they still ignore many of the details of rotational spin-down physics, such as changes in the magnetic field properties and wind acceleration.

3.1.2 Radio observations

Several attempts have been made to observe stellar winds at radio wavelengths, and while such winds have not been directly detected, this technique has been able to put important upper limits on mass loss rates. These upper limits can be derived in two ways. Given their temperatures and ionization states, it is expected that free-free thermal radiation at radio wavelengths will be detectable if the winds are sufficiently luminous ([153];[210]). The non-detection of radio emission from a star's wind can therefore be used to derive an upper limit on the wind's density and mass loss rate. For example, [62] was unable to detect emission from the winds of three young Sun-like stars and found that the mass loss rates must be below $\sim 5 \times 10^{-11} M_{\odot} \text{ yr}^{-1}$. Possibly even more useful is the fact that with a sufficiently dense wind, radiation emitted from a star's surface at some radio wavelengths will not escape the system. The observations of flares on the surfaces of stars in these wavelength regions imply that the winds are not sufficiently dense, which can be used to derive upper limits on mass loss. This method was used by [121] to derive an upper limit on the mass loss of the $0.3 M_{\odot}$ M-dwarf YZ CMi of $\sim 10^{-12} M_{\odot} \text{ yr}^{-1}$. Note that in some cases, stellar winds can also cause radio emission from the magnetospheres of planets with short orbital periods to be unobservable ([195]). The most sensitive upper limits on the winds of young solar mass stars were derived by [56] who found that the mass loss rates of the 300 Myr old π^1 UMa and the 650 Myr old κ^1 Ceti are below $\sim 5 \times 10^{-12} M_{\odot} \text{ yr}^{-1}$. These upper limits suggest that very active stars likely have mass loss rates that are within a factor of 100 of that of the current Sun, though they say little about the winds of very young pre-main-sequence stars.

3.1.3 Interactions with the interstellar medium

It is possible that wind properties can be inferred from their interactions with the interstellar medium (ISM) and stellar astrospheres, which are the stellar equivalents of the heliosphere. At distances of several tens of AU from the Sun, the solar wind begins to interact strongly with neutral hydrogen from the ISM. Protons in the solar wind charge exchange with ISM hydrogen creating a slowly moving ion and a neutral hydrogen atom traveling at several hundred km s^{-1} called an energetic neutral atom (ENA). Similar charge exchange reactions involving heavy ions in the wind lead to the emission of X-ray photons and [200] attempted to observe these photons in the astrosphere of Proxima Centauri. Their lack of detection implied an upper limit on the mass loss rate of $\sim 3 \times 10^{-13} M_{\odot} \text{ yr}^{-1}$.

The slowly moving ion created in charge exchange reactions is picked up by the solar wind and this has the effect of slowing the wind and eventually making it subsonic. If enough neutral hydrogen is present in the region of the ISM that a stellar system is embedded in, a wall of neutral hydrogen can build up at the edge of the astrosphere ([207]). If our viewing angle to the star is such that the star's radiation passes through this neutral hydrogen wall, much of the star's Ly- α emission line will be absorbed. Unfortunately, due to absorption by neutral hydrogen in the ISM, only a small fraction of a star's Ly- α emission is observable, which makes detecting the signature of this astrospheric absorption challenging. The results can be compared to models of wind-ISM interactions and used to estimate wind properties. [207], [209], [206], [208] used this technique to estimate wind properties of several stars and found that stars with stronger X-ray emission have higher mass loss rates. For example, they estimated a mass loss rate for the $0.82 M_{\odot}$ K dwarf ϵ Eri to be $\sim 6 \times 10^{-13} M_{\odot} \text{ yr}^{-1}$, which is

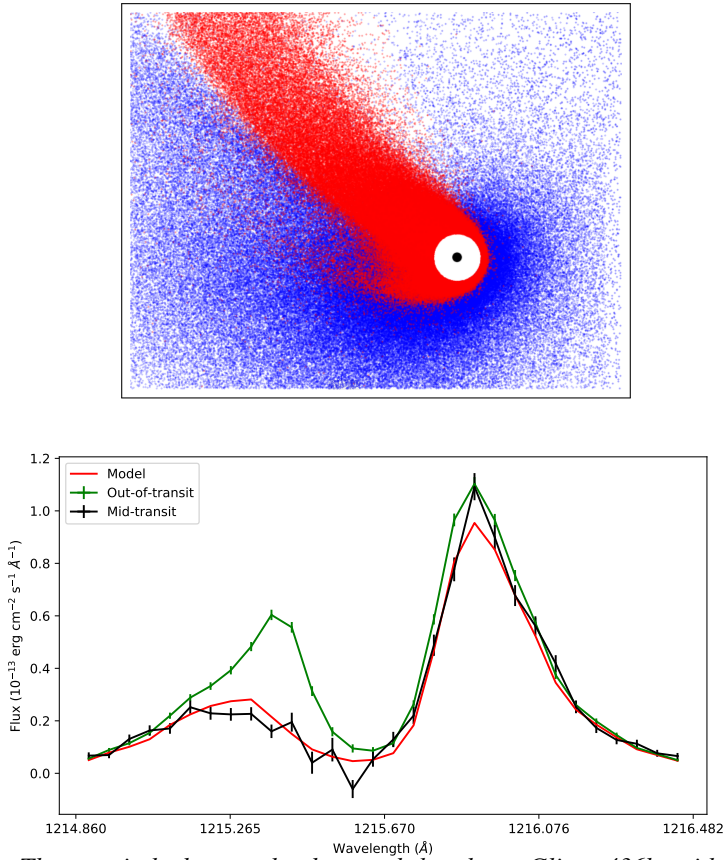


Figure 80: *The atomic hydrogen cloud around the planet Gliese 436b, with red and blue showing neutral and ionised hydrogen atoms (each dot corresponds to 1×10^{32} atoms) and the black circle showing the planet's size. The lower panel shows the observed Ly- α emission line of the star in and out of transit and the mid-transit model simulated by [100]. Images courtesy of K. G. Kislyakova.*

a factor of 30 above that of the current Sun. Their results suggested that $\dot{M}_w \propto F_X^{1.34}$, which for the solar wind at later ages suggests $\dot{M}_w \propto t^{-2.33}$.

The derived relation between X-ray activity and mass loss breaks down for the most active stars which all show much weaker than expected mass loss. For example, their measured mass loss from the young 300 Myr old solar analogue π^1 UMa was half that of the modern solar wind ([208]). These results suggest that very active stars could have much weaker winds than we would expect, and could mean that for the first Gyr of the solar system's history, the solar wind mass flux was similar to modern values. However, as discussed in [90], this interpretation is difficult to reconcile with our empirical understanding of stellar rotational evolution. The rapid spin-down of rapidly rotating active stars that we infer from young stellar clusters requires high wind mass loss rates. The dependence of the solar wind mass loss rate on $t^{-2.33}$ at later ages is also much stronger than the results implied by rotational evolution models which suggest a t^{-1} dependence.

3.1.4 Effects on planetary companions

Several techniques have been suggested for inferring wind properties indirectly using measurements of the influences of these winds on transiting planets. Planets are surrounded by a large and tenuous cloud of particles originating from their upper atmospheres called the exosphere and neutral particles in the exosphere charge exchange with stellar wind protons creating an tail of ENAs flowing away from the planet, as shown in Fig. 80. While charge exchange with heavy ions in the wind produces X-ray photons, this emission is unobservable outside our solar system ([99]). Neutral hydrogen atoms in both the exosphere and in the ENA tail can transit the star and the resulting absorption causing variability of the star's Ly- α emission line over the course of each orbit. This can be seen for example in the Ly- α line of the M dwarf Gliese 436 which is orbited by a warm Neptune-sized transiting planet as shown in Fig. 80. This absorption has been observed for several transiting planets with hydrogen dominated primordial atmospheres ([192]; [49], [50]).

The time variability of the Ly- α line for a given star-planet system depends on the properties of the star's wind in the vicinity of the planet and it is possible to use this to derive knowledge about the winds of such stars. [105] compared the Ly- α line of HD 209458, which is transited by a short period planet with a mass slightly below that of Jupiter, to models for interactions between the star's wind and the planet's exosphere. They were able to estimate a wind speed in the region of the planet of 400 km s^{-1} and they were also able to estimate the magnetic moment of the planet which is important because measurements of exoplanet magnetic moments are very challenging ([203]). Similarly, [20] derived a stellar wind speed in the vicinity of the transiting planet GJ 436b of 85 km s^{-1} and [193] used their constraints to model the host star's wind, finding a mass loss rate of $\sim 10^{-15} \text{ M}_{\odot} \text{ yr}^{-1}$. Other methods for constraining stellar wind properties using transiting exoplanets have been explored ([124]).

Another promising avenue for measuring stellar wind properties involves hydrogen atmosphere white dwarfs that have spectral absorption lines from metals, which in many cases is unexpected given that metals should rapidly diffuse out of the upper atmosphere. It is possible that in many systems, these metals could come from the winds of companion M dwarfs accreted onto the surfaces of the white dwarfs ([156]). [43] explored the possibility of constraining the mass loss of the M dwarf companions of six such white dwarfs, three of which have small orbital separations of 0.015 AU or less and the other three have separations of 1.5–159 AU. For the close orbit systems, they estimated mass loss rates between 10^{-16} and $6 \times 10^{-15} \text{ M}_{\odot} \text{ yr}^{-1}$ and for the wider systems, they estimated values above $10^{-10} \text{ M}_{\odot} \text{ yr}^{-1}$, which they argue are unrealistically large.

3.2 Solar wind: basic properties

Although a simplification, it is useful here to describe the solar wind as being composed of three components: these are the slow wind, the fast wind, and coronal mass ejections (CMEs). Typically, the solar wind is not isotropic but has variable speed in different directions with speeds typically between 300 and 900 km s^{-1} and the distribution of wind speeds shows two clear components with average speeds of 400 km s^{-1} for the slow component and 760 km s^{-1} for the fast component. The slow wind is cooler and more variable on short timescales than the fast wind and the two components have different abundances of heavy ions ([197]). However, the two components have very similar mass fluxes, meaning that the outflow mass flux from the Sun is approximately isotropic and the mass loss rate is approximately constant over the solar cycle.

If we do not consider transitory events like CMEs (discussed instead in Section 3.5), the main change in the solar wind properties over the solar cycle is the spatial distribution of

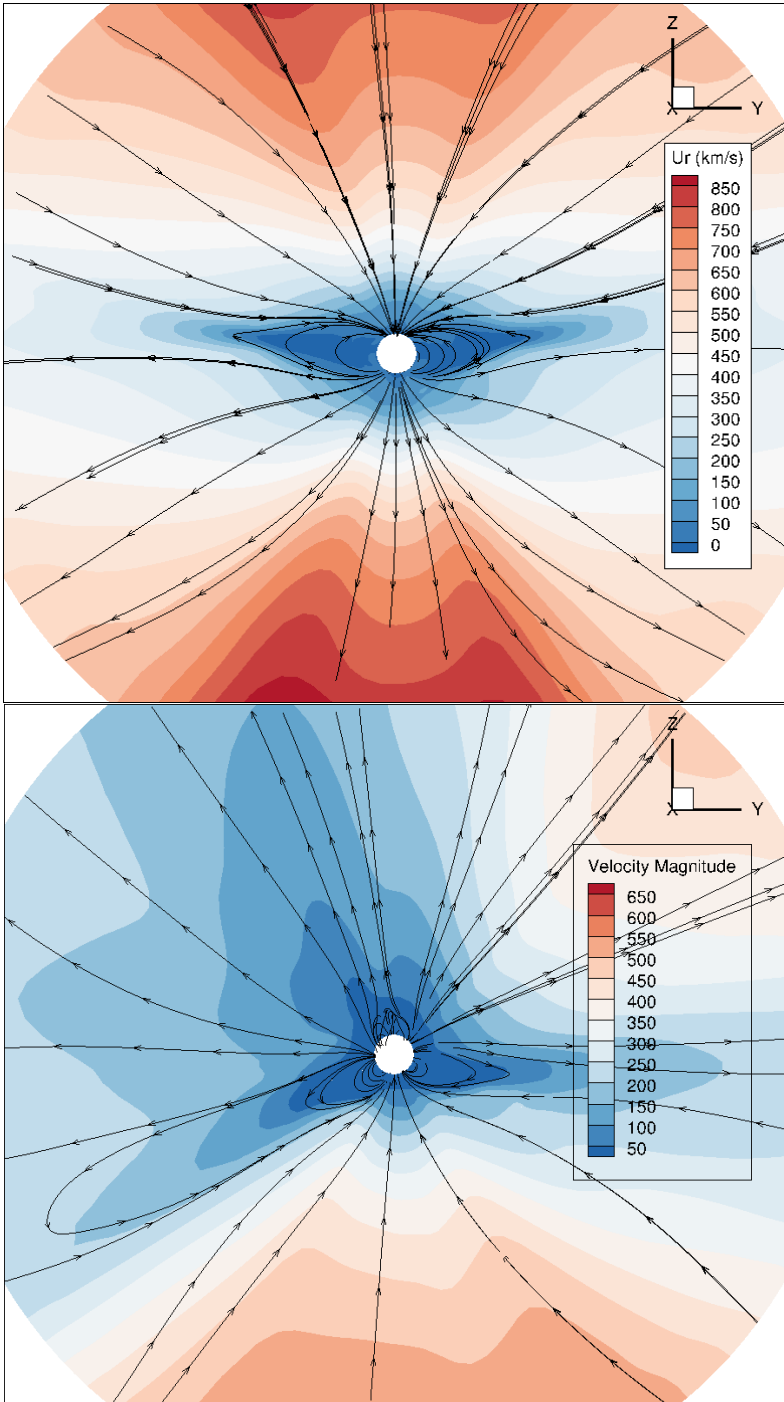


Figure 81: Magnetic field and velocity structure in the solar wind with dipolar (upper panel) and more complex (lower panel) field structures as simulated by [18] using the Space Weather Modelling Framework ([187]). Black lines show magnetic field lines and colored contours show wind speed and the close relation between acceleration and magnetic field can be seen. Images courtesy of S. Boro-Saikia.

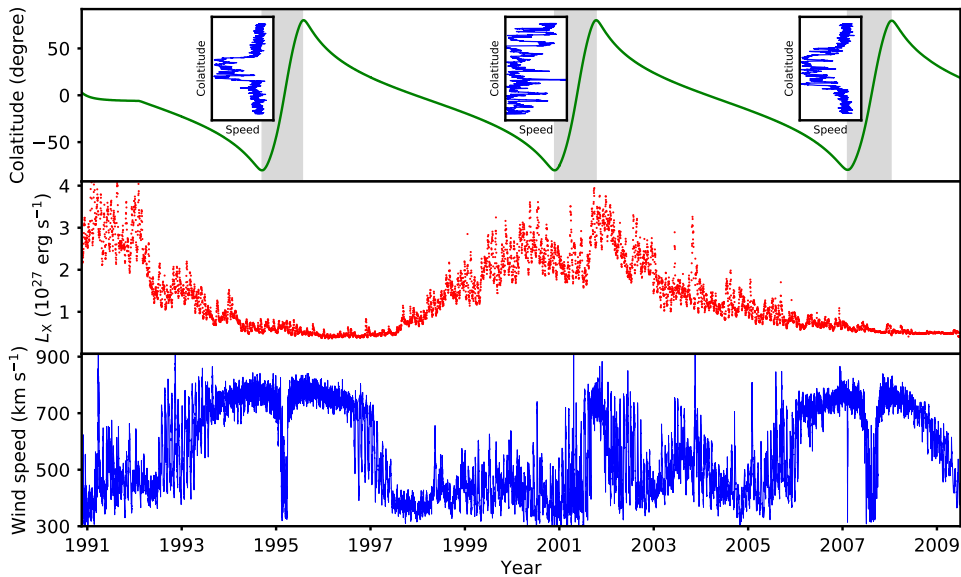


Figure 82: Solar wind speed as a function of latitude and solar cycle as revealed by the Ulysses spacecraft between 1990 and 2010, with Ulysses measurements obtained from <http://cohoweb.gsfc.nasa.gov>. The upper panel shows the colatitude of Ulysses and the shaded areas indicate the three fast latitude scans. The insets next to each fast latitude scan shows the wind speed measured by the spacecraft as a function of colatitude for the three scans and we can easily see the differences in wind structure between cycle minimum (first and third scans) when the Sun's global magnetic field was dipolar and maximum (second scan) when the field was more complex. The middle panel shows solar X-ray luminosity during this time, with the L_X values derived from the Flare Irradiance Spectrum Model ([24]) and the solar cycle variability is clearly visible. The lower panel shows wind speed measured by Ulysses during the time.

slow and fast wind. These two components are closely related to the geometry of the Sun's global magnetic field, which is typically very simple and dipolar at activity minimum and much more complex and multipolar at activity maximum. As can be seen in Fig. 81, we can break down the Sun's global magnetic field into regions of closed and open field lines and it is known empirically that fast wind originates from regions of open magnetic field, whereas the slow wind originates from above closed magnetic field ([198]). However, despite this clear correlation, the exact physical origin on the Sun's surface of the slow component is unclear and various possibilities, such as the boundaries between open and closed field lines and active regions within closed field regions, have been suggested ([3]). At solar minimum, fast wind originates from high latitudes and slow wind dominates in equatorial regions, whereas at solar maximum, the distribution of fast and slow wind is much more complex and unpredictable. This can be seen from the wind speed measurements of the Ulysses spacecraft which had a very high orbital inclination relative to the plane of the solar system and was therefore able to measure the solar wind properties at all latitudes. Wind speed measurements and the relation to latitude and the solar cycle from Ulysses are shown in Fig. 82.

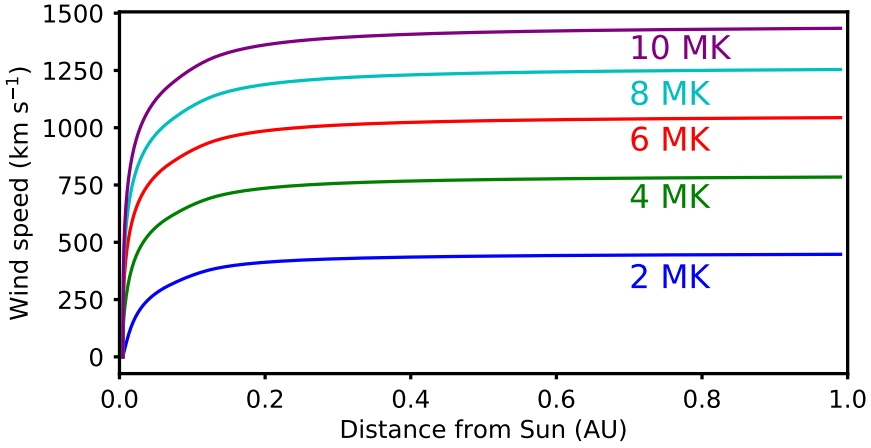


Figure 83: Wind speed as a function of distance from the center of the Sun for different lower boundary temperatures assuming a thermal pressure driven wind only and a polytropic equation of state as calculated using the model of [92].

3.3 Wind physics

After the initial suggestion by [17] of the presence of a fast flowing wind propagating outwards from the Sun, [155] derived a simple hydrodynamic model for the radial outflow of the solar wind. In this model, the wind is accelerate away from the Sun by thermal pressure gradients, which is possible given that the outer atmosphere of the Sun reaches temperatures of million of K and the assumption by [155] that the wind is isothermal. This model was extended to take into account the effects of the Sun's rotation and magnetic field on the wind by [204] who derived for the radial acceleration of the wind

$$V_r \frac{dV_r}{dr} = -\frac{1}{\rho} \frac{dp}{dr} - \frac{GM_\star}{r^2} + \frac{V_\phi^2}{r} - \frac{B_\phi}{4\pi\rho r} \frac{d}{dr}(rB_\phi), \quad (62)$$

where r is the distance from the center of the star, V_r is the radial outflow speed, ρ is the mass density, p is the thermal pressure, and V_ϕ and B_ϕ are the azimuthal components of the wind velocity and magnetic field. The four terms on the right hand side of this equation correspond to the acceleration due to thermal pressure gradients, gravity, the centrifugal force, and the radial component of the Lorentz force. Due to the Sun's slow rotation and relatively weak magnetic field, the third and forth terms are negligible for the solar wind and when ignored, this equation gives the radial acceleration from the model by [155]. In this model, the wind speed far from the surface of the star is determined primarily by temperature, with higher temperature winds having higher outflow speeds. The mass flux in the wind is determined by both the temperature and the density at the base, wherever that is defined. The isothermal assumption leads to unrealistically large acceleration far from the Sun and a slightly more realistic model can be obtained by assuming a polytropic equation of state, meaning that $p \propto \rho^\alpha$, where $\alpha = 1$ and $\alpha = 5/3$ correspond to isothermal and adiabatic winds respectively (though a Parker wind cannot form if the gas is adiabatic). Note that this is not the relation $p \propto \rho^\gamma$ for an adiabatic gas and α is not the adiabatic index. Within the solar corona, α is approximately 1.05 ([182]) and in the inner heliosphere, α is approximately 1.46 ([188]). Wind acceleration profiles for a polytropic wind with different base temperatures are shown in Fig. 83 and a detailed explanation of these types of models is given by [114].

The main problem with the models described above is the lack of any description of how the wind is heated, with the temperature and density at the base of the wind being free parameters. Since we have no clearly reliable method for determining these parameters for other stars, a primary goal for solar and stellar wind physics is the identification and description of the mechanisms responsible for heating the outer atmospheres and winds of stars. It is generally accepted that the source of the energy is convective motions in the photosphere and this energy is transferred to and released in the corona and wind by magnetic fields, but the mechanisms responsible are currently disputed. As described in detail in [37], physical wind heating models break down into two classes: these are wave/turbulence driven models and reconnection/loop-opening models. The former involve waves created by the jostling of magnetic field lines in the photosphere propagating upwards and being dissipated at higher altitudes. These are typically assumed to be Alfvén waves and many physical models that heat winds in this way have been developed ([40]; [183]; [190]). The latter involve the dragging around of magnetic field lines by photospheric convective motions causing the field lines to become stressed, building up energy in the magnetic fields that is then released by reconnection. Currently Alfvén wave heating is the favored explanation and is used as the main wind driving mechanism in 3D global wind models ([191];[164]). For a recent review, see [41].

A further complication to the picture involves measurements of the solar wind temperatures near and far from the Sun. Temperatures derived from the Solar Heliospheric Observatory (SOHO) mission have shown that the gas within coronal holes is typically between 1 and 2 MK ([55]; [9]) and if we assume an isothermal Parker wind, these temperatures are high enough to accelerate the wind to speeds typical for the fast solar wind. However, *in situ* measurements of the wind far from the Sun show that it is not isothermal and the hotter fast component typically has temperatures of $\sim 3 \times 10^5$ K at 1 AU. While this is much hotter than we would expect if the solar wind was cooling adiabatically as it expands, meaning that wind heating must take place throughout the inner heliosphere, models that correctly reproduce the temperature structure of the wind are unable to explain the outflow speeds of the fast wind using thermal pressure gradients as the only acceleration mechanism (e.g. see Fig. 4 of [92]). It is therefore necessary that another acceleration mechanism not included in Eqn. 62 is important, and a prime candidate for this is momentum deposition from waves. [35] estimated that close to the solar surface, thermal pressure gradients dominate the acceleration, and by a few solar radii, direct momentum dissipation from waves contribute approximately half of the acceleration (see their Fig. 3).

3.4 Winds of active stars

As discussed earlier, little is known about the properties of the winds of active stars largely due to our lack of a theoretical understanding of the underlying mechanisms that heat and accelerate the solar wind and our lack of observational constraints on wind properties. While many models have been developed to predict the properties of active star winds, we have no clear way to assess the validity of the assumptions made or to test the predictions of the models. In fact, it is reasonable to wonder if we know anything about the winds of active stars other than that they are likely to have outflow speeds that are within a factor of a few of the solar wind speed and mass loss rates that are within a few orders of magnitude of the modern Sun's mass loss rate. Given the observed rotational evolution of rapidly rotating stars, we have good reason to believe that more active stars have higher mass loss rates and this saturates at approximately the same rotation rate that X-ray emission saturates. A plausible scenario for wind mass loss as a function of rotation is shown in Fig. 84, where the increase in loss rate for very rapid rotation is discussed below.

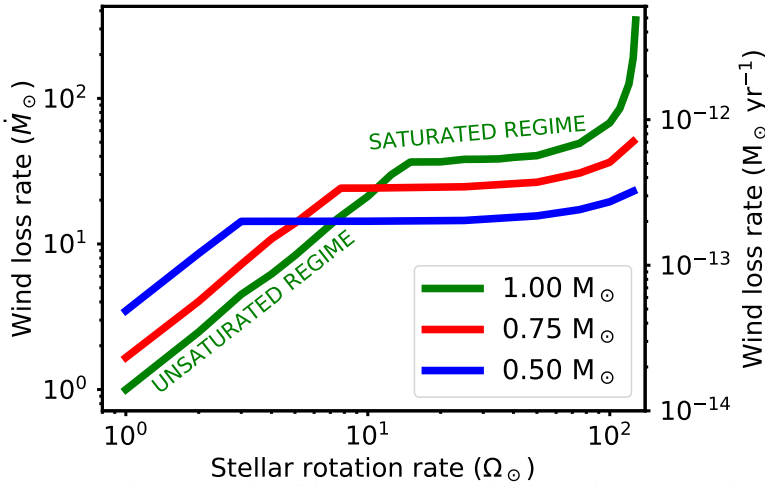


Figure 84: Figure showing a plausible scenario for the dependence of wind mass loss on rotation for several stellar masses based on the 1D MHD wind model described by [87]. The mass loss rates for solar mass stars increase rapidly on the right side of the figure as the line approaches the break-up rotation rate.

The coronae of more active stars are hotter than those of less active stars. Since both the Sun's X-ray emitting corona and the solar wind are heated by energy transferred from the photosphere by the magnetic field, we might expect that more active stars have hotter winds. This is plausible and would mean that the wind speeds and mass loss rates are higher for active stars and decay on the main-sequence due to rotational spin-down ([80]; [90]). However, while the link between the X-ray emitting coronae and winds of active stars is plausible, several issues exist. It could be that the mechanisms responsible for heating winds and coronae are different ([39]); for example, it could be that the corona on small scales is heated by magnetic reconnection while the wind is heated by the dissipation of Alfvén waves. If true, these different mechanisms could respond to changes in a star's magnetic field differently. The approximately constant mass loss rate of the Sun over the cycle despite large changes in the X-ray luminosity makes the link between coronal activity and winds unclear ([30]) though we should be cautious with interpreting this lack of correlation since the spin-down related decay in stellar magnetic fields and X-ray emission is not analogous to the changes in the Sun's magnetic field and X-ray emission over the solar cycle. Even if the winds of more active stars are hotter, [183] suggested that the higher densities in the wind cause enhanced radiative cooling that can reduce the final speeds and saturate the mass loss.

For very rapidly rotating stars, other processes become important. Although the third and forth terms on the right hand side of Eqn. 62 for the centrifugal and Lorentz forces respectively are negligible for the Sun, they are significant for the winds of more strongly magnetized and more rapidly rotating stars. This was studied by [13] who identified two distinct regimes. Stars are in the slow magnetic rotator regime when centrifugal and Lorentz forces have a negligible contribution to the acceleration of the wind. Stars are in the fast magnetic rotator regime when these forces dominate in determining the flow speeds of the wind far from the star. In this regime, the mass loss rate is still determined primarily by other processes. It is also possible for very rapidly rotating stars to be in a third regime, called the centrifugal magnetic rotator regime, where centrifugal and Lorentz forces influence significantly the wind mass loss rates. A star enters this regime when the equatorial corotation radius is below the sonic point. The properties of winds for stars that are very rapidly rotating

were studied by [87] who showed that as stars approach the break-up rotation rate, wind mass loss rates rapidly increase as shown in Fig. 84, though they also found that the wind properties in the fast and centrifugal rotator regimes depend on assumptions made for the wind temperatures of active stars.

3.5 Coronal mass ejections

There is another factor that complicates the picture and adds another level of uncertainty to our understanding of the winds of active stars. Coronal mass ejections (CMEs) are eruptive events involving a bulk of plasma with its magnetic field that is accelerated away from the Sun and propagates through interplanetary space. Solar CMEs have a very large range of properties and typically have speeds between 100 and 1000 km s⁻¹ and masses between 10¹³ and 5 × 10¹⁶ g ([2]). Their angular widths can be as small as 10° or larger than 180° ([214]). The rate at which the Sun produces CMEs varies with the solar cycle, with a higher CME occurrence rate being seen at solar maximum ([202]). There is a clear link between CMEs and solar flares ([75]; [2]) and it is likely that in many cases, flares and CMEs result from the same magnetic events, but the exact nature and universality of this link are unclear. While flares and CMEs often occur together, many CMEs are observed without a corresponding flare and many flares are observed without a corresponding CME, though this is surely influenced by difficulties in detecting and linking these two phenomena. It appears that the link with flares is stronger for larger solar flares, and [199] estimated that 90% of X-class flares (the largest class of flares) are associated with CMEs. More energetic flares are associated with more massive and rapidly outflowing CMEs. For a review of solar CMEs, see [28].

Since more active stars have higher flare occurrence rates, we should expect that they have also higher CME rates. This is difficult to study since, like winds, CMEs are very difficult to detect observationally on other stars ([117]) though some possible detections exist ([140]; [11]; [118]). One way to study CMEs from active stars is to combine statistical relations for the solar flare–CME relation with observationally determined statistics for flare rates and energies on active stars. [1] studied CME winds for pre-main-sequence stars and estimates mass loss rates of between 10⁻¹² and 10⁻⁹ M_⊙ yr⁻¹. For saturated Sun-like main-sequence stars, [45] estimated CME mass loss rates of 10⁻¹⁰ M_⊙ yr⁻¹. For a review of these types of extrapolations, see [147]. These high CME rates would lead to significant angular momentum removal and could be one of the factors influencing the rotational evolution of rapidly rotating stars ([170]; [1]; [85]).

These results suggest that the winds of very active stars could be entirely dominated by CMEs and planets orbiting such stars could experience highly turbulent space weather conditions. However, several issues exist with these estimates. The high mass loss estimated by [45] leads to a continuous CME energy loss that is 10% of the star's bolometric luminosity which is unreasonably large. They suggested that the flare–CME relation must be invalid for the largest flares to avoid such extreme energy loss. A possible solution to this problem is the suppression of stellar CMEs by the global magnetic fields of host stars ([6]). While the global component of the Sun's magnetic field is typically ~1 G, young active Sun-like stars have values that are orders of magnitude larger ([58]) and these very strong fields can trap CMEs within the closed corona.

4 Earth's Upper Atmosphere and Wind Interactions

To understand the diverse ways that stellar winds can influence the atmospheres of planets, it is useful to consider first the modern Earth's since this is the most well understood case.

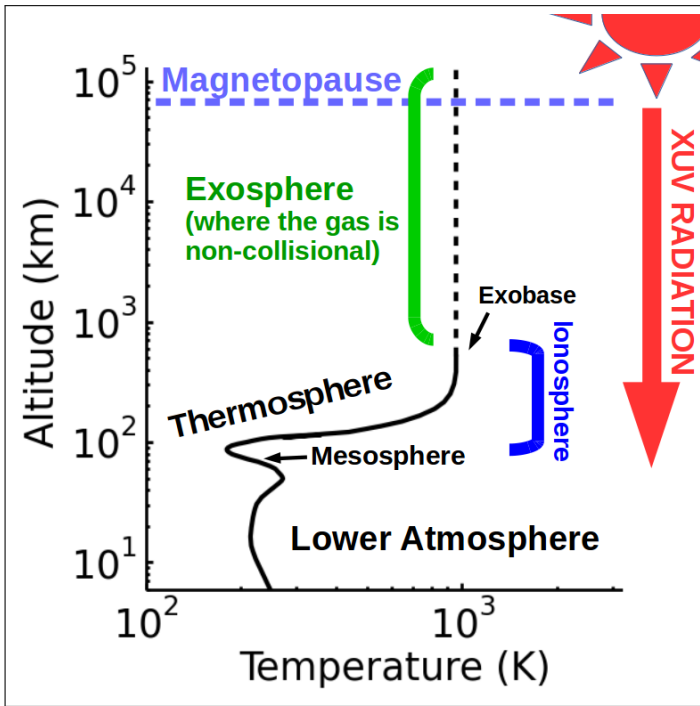


Figure 85: Cartoon showing the vertical structure of Earth's atmosphere.

The modern Earth's atmosphere is composed primarily of N_2 and O_2 (and 1% Ar which is of little importance), with small amounts of other species such as CO_2 and H_2O , and the Sun's magnetic activity is generally very low. In this section, I describe the structure and physics of the Earth's upper atmosphere and magnetosphere, and in the next section, I describe how this is likely to be different for planets with different types of atmospheres orbiting different types of stars.

4.1 The vertical structure of the Earth's atmosphere

It is common to break down the vertical structure of the atmosphere into layers based on the temperature gradient as shown in Fig. 85. The lowest layer is the troposphere which extends from the surface to the tropopause at an altitude of 10 km and has a decreasing temperature with increasing altitude (i.e. a negative temperature gradient). The troposphere is heated by its contact with the surface and this heat is distributed upwards by convection. Above the troposphere is the stratosphere, which has a positive temperature gradient due to the absorption of solar ultraviolet radiation by ozone and extends from 10 km to the stratopause at approximately 50 km. These two regions are labeled in Fig. 85 as the 'lower atmosphere'. Above this is the mesosphere, which again has a negative temperature gradient due to cooling by the emission of infrared radiation to space by CO_2 molecules and extends from 50 km to the mesopause at approximately 100 km. Above this is the thermosphere, which has a strong positive temperature gradient due to the absorption of solar X-ray and extreme and far ultraviolet radiation, particularly by O_2 and O . The temperature of the Earth's thermosphere typically varies between approximately 500 and 1500 K and depends primarily on the Sun's activity ([166]). The thermosphere extends to the exobase, which is typically at an altitude

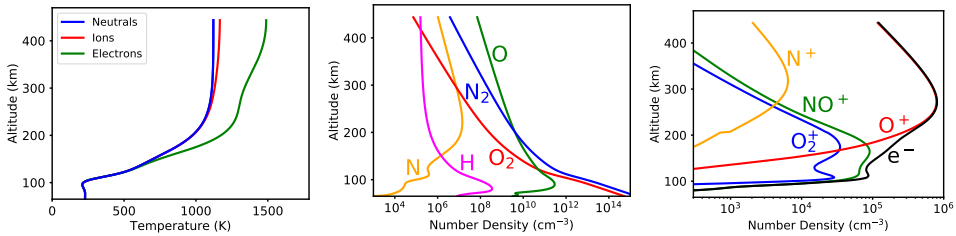


Figure 86: *The temperature and chemical structure of the Earth's upper atmosphere as calculated by [91].*

of a few hundred km and varies also with the Sun's activity. The exobase is where the gas density drops low enough that the gas is approximately non-collisional.

It is useful also to consider the chemical structure of the atmosphere which also has several layers, the most important of which are the homosphere and the heterosphere. The homosphere extends from the surface until approximately 120 km and in this region the gas is well mixed by turbulent processes, causing the mixing ratios of stable long-lived species (e.g. N_2 , O_2 , CO_2 , Ar, etc.) to be uniform with altitude regardless of their masses. An important exception to this is water vapor which has a very low abundance at high altitudes due to the presence of the cold trap in the lower stratosphere which prevents water molecules from reaching the upper atmosphere. The upper boundary of the homosphere is called the homopause or turbopause. Above the homopause is the heterosphere and in this region molecular diffusion separates the chemical species by mass such that heavier species become increasingly less abundant at higher altitudes. The Sun's XUV radiation dissociates molecules and causes the gas throughout most of the thermosphere to be dominated by atomic O and N as shown in Fig. 86.

If we consider a hydrostatic atmosphere only, the density of the i th chemical species, n_i , decreases with altitude, z , as $n_i \propto \exp(-z/H_i)$, where H_i is the pressure scale height given by kT/mg , where k is the Boltzmann constant, T is temperature, m is the molecular mass of the gas, and g is the gravitational acceleration. Note that this only applies to long lived species that are not rapidly created and destroyed by chemical processes. The difference between the homosphere and the heterosphere is the molecular mass term, m . In the homosphere, this is the average molecular mass of the entire gas meaning that each chemical species has the same scale height regardless of mass. In the heterosphere, this is the mass of individual species, meaning that heavier species have smaller scale heights.

The upper mesosphere and thermosphere correspond to the ionosphere where the absorption of solar X-ray and EUV photons causes ionisation of the gas. The density of ions depends sensitively on the Sun's activity and varies over the solar cycle (e.g. [178]). For the modern Earth, the ionisation fraction of the gas reaches values of $\sim 10^{-3}$ at typical solar maximum conditions and this small mixing of ions is very important for interactions between the upper atmosphere and the magnetosphere. By definition the ionosphere extends to the exobase and since the gas densities are low in the upper thermosphere, the neutral, ion, and electron components of the gas have different temperatures. The thermal and chemical structure of the ionosphere is shown in Fig. 86.

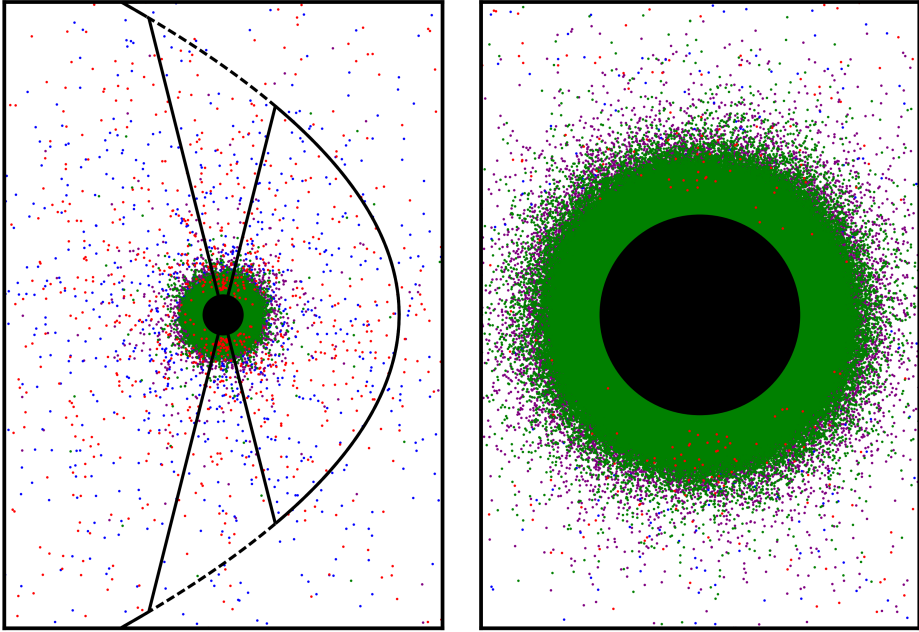


Figure 87: *The basic structure of the exosphere of the Earth assuming the Sun's activity from 3 Gyr ago as calculated using a kinetic particle model by [102], with each point corresponding to approximately 10^{29} particles. The right panel shows the same as the left panel but zoomed in around the planet which is shown in black. The purple and green points show atomic nitrogen and oxygen neutrals respectively, and the blue and red points show atomic nitrogen and oxygen ions. The black parabola in the left panel shows the approximate boundaries of the magnetosphere and the wedge shaped lines coming from the poles show the polar opening angle expected 3 Gyr ago due to stronger solar wind conditions. Image courtesy of K. G. Kisylakova.*

4.2 Exospheres, magnetospheres, and winds

Above the exobase is the exosphere, which is approximately isothermal and extends out into interplanetary space (there is no clear and exact definition for the upper boundary of the exosphere). A useful definition for the exobase is based on the Knudsen number, Kn , which is a dimensionless quantity given by the ratio of the mean free path of particles to a relevant length that characterises the scale of the system, which for planetary atmospheres is typically the pressure scale height, which we can define as $H = -p/(dp/dz)$, where p is the thermal pressure and z is the altitude or radius, and for a fully hydrostatic atmosphere this is given by kT/mg . Low in the Earth's atmosphere, the high density of the gas means $Kn \ll 1$ and as we go to higher altitudes this value increases until the exobase where it becomes unity. Below the exobase, the particle speed distribution can be approximated by a normal thermal Maxwellian distribution. In the exosphere, the much more tenuous gas means that $Kn \gg 1$ and particles are mostly able to move freely with few collisions and they do not follow the Maxwellian distribution. A particle simulation of the exosphere of the Earth, assuming solar conditions from 3 Gyr ago, is shown in Fig. 87.

Particles that cross the exobase can have a range of fates depending on their velocities and their interactions with other exospheric particles ([16]) and with the solar radiation field

and wind ([216]; [15]). In the absence of any interactions, particles either fall back to the atmosphere on ballistic trajectories or escape the system entirely depending on if they are traveling slower or faster than the escape velocity. If they collide with other exospheric particles, they can become satellite particles that orbit the planet. Particles can also become ionised by absorbing solar XUV photons, by colliding with electrons in the solar wind, or by charge exchanging with ions in the solar wind. If ionised, their trajectories are also influenced significantly by the Earth's magnetosphere or by the solar wind depending on where the ionisation takes place, with ions created beyond the boundary of the magnetosphere being swept away by the solar wind magnetic field. As discussed in Section 3.1.4, charge exchange between exospheric particles and wind protons creates a population of fast moving neutrals called ENAs. Since these ENAs are not influenced by the planet's magnetosphere, ENAs created between the planet and the direction from which the wind is propagating enter the atmosphere and collide with atmospheric particles, which can heat the upper atmosphere ([26]; [119]).

When considering atmospheric loss processes, an important source of ions in the magnetosphere is upward flows from the ionosphere. Given their much lighter mass, ionospheric free electrons are poorly bound gravitationally and expand away from the planet causing a reduction in the electron density and the generation of an ambipolar electric field. This field acts to maintain quasi-neutrality in the gas by inhibiting the outflow of electrons, which causes an outward pressure on the ions resulting in a flow of ions and electrons from the ionosphere into the magnetosphere. These particles are then distributed throughout the magnetosphere ([47]) and their ultimate fates are currently a matter of debate. While some studies suggest that most escape the magnetosphere and are lost to interplanetary space ([145]), others suggest that most are unable to escape and are recaptured by the planet ([172]). The outward flow of ions and the fraction of these ions that escape to space depend on the solar wind conditions and emission of high-energy radiation ([74]). For a review of ion flows in the magnetosphere, see [138].

The structure of Earth's magnetosphere is shown in Fig. 88. For the Earth, the boundaries of the magnetosphere are very large in comparison to the planet and the atmosphere. The outer boundary of the magnetosphere is called the magnetopause and in the direction of the incoming solar wind is approximately at the point where the magnetic pressure from the Earth's magnetic field balances the ram pressure in the wind, which is typically at a radius of $\sim 10 R_{\oplus}$. Since the exobase is typically at a radius of $1.07 R_{\oplus}$, the very large magnetosphere compared to the thickness of the atmosphere means that most exospheric particles do not interact directly with the solar wind. For this reason, the magnetosphere is often described as protecting the atmosphere from escape to space by wind erosion. However, the situation is more complicated and there is disagreement about the effects of the magnetosphere on escape, as discussed below. Note that planets that do not possess significant intrinsic magnetic fields, such as Venus and Mars, form induced magnetospheres due to interactions between the solar wind and their ionospheres and exospheres ([14]). The interaction between the solar wind and the Earth's magnetosphere also causes additional effects lower in the atmosphere, especially during geomagnetic storms when the rapid changes in the space weather conditions in Earth's vicinity causes reconnection in the magnetosphere and jostling of the magnetic field lines that connect to the Earth's ionosphere ([25]). The former causes the acceleration of particles that can precipitated downwards into the thermosphere and below, driving thermal and chemical processes. The latter causes electric currents in the ionosphere, and as these charged particles collide with neutrals, energy is dissipated causing the gas to be heated (this process is called Joule heating), especially in the region above 100 km ([166]).

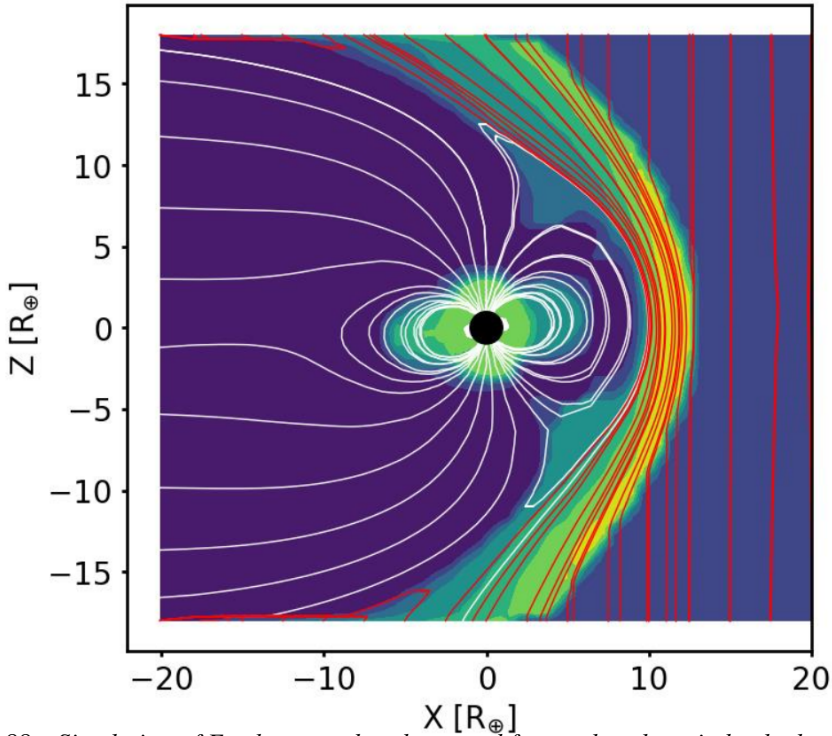


Figure 88: *Simulation of Earth exposed to dense and fast early solar wind calculated using the Space Weather Modelling Framework ([187]). The background color shows the plasma density, and red and white lines illustrate the planetary and solar wind magnetic field lines. Image courtesy of K. G. Kislyakova.*

5 Planetary Upper Atmospheres and Escape

Loss processes take place in the planet's thermosphere and exosphere and are mostly driven by the central star's output of high-energy radiation and wind. These processes are often broken down into thermal and non-thermal processes (Lammer 2013) where the thermal processes are Jeans escape and hydrodynamic escape and non-thermal processes include everything else. Jeans escape involves the high energy tail of the thermal Maxwellian distribution at the exobase with upward speeds greater than the planet's escape velocity, allowing the particles to leave the planet entirely. This is especially important for low mass particles such as hydrogen, but can also be important for higher mass particles if the atmosphere is very hot. Several processes in the planet's exosphere ionize neutral particles leading them to be picked up by the stellar wind ([101]; [20]). Other significant processes include the outflow of ions from the magnetic poles as discussed above ([68]) and photochemical escape, which involves the escape high speed particles created in chemical reactions ([8]) and is powered by energy absorbed from the star's XUV spectrum. In all these cases, the loss rates are influenced heavily by the chemical and thermal structure of the upper atmosphere, with hotter atmospheres having higher loss rates, in some cases because they are more expanded and therefore more exposed to direct interactions with the star's wind. When gas temperatures are very high, it is possible for the upper atmosphere to flow away from the planet hydrodynamically (e.g. [186]) at very high rates. The structure of the upper atmosphere is determined

by many thermal and chemical processes and is influenced by several factors, including the mass of the planet, the chemical composition of the atmosphere, and the star’s X-ray and ultraviolet spectrum.

It is important to understand what effect an intrinsic planetary magnetic field has on these escape processes and on the total escape. It is often assumed that by reducing the direct interactions between a planet’s exosphere and the star’s wind, magnetic fields protect atmospheres from escape. However, there are a large number of loss processes that respond to planetary magnetic fields in complex ways and the net effect of a magnetosphere is not clear. While it is true that direct ionisation and pickup of exospheric neutral particles by winds is prevented by magnetospheric shielding, the magnetosphere collects energy from the wind and dissipates this energy lower in the atmosphere, driving processes such as high energy particle acceleration and Joule heating. [73] studied the importance of the presence and strength of a planet’s intrinsic magnetic field on several loss processes and found that in some cases, the magnetic field increases the escape of ions. It is possible that given the complexity of the problem, a planetary magnetic field influences escape in different ways for different parts of the parameter space. For example, this was found by [48] who studied ion escape from weakly-ionized Mars-sized planets.

[73] pointed out that the magnetised Earth and the unmagnetised Venus and Mars all lose atmosphere at approximately the same rate (~ 0.5 to 2 kg s^{-1}). While this is suggestive, differences in the masses, orbital distances, and atmospheric compositions of these planets possibly explain the similar loss rates. For example, while Venus is closer to the Sun and lower mass than Earth, its atmosphere is composed mostly of carbon dioxide which, as I explain below, cools the upper atmosphere and protects it from escape, whereas Earth’s N_2 and O_2 dominated atmosphere heats up and expands much more readily. Mars is further from the Sun than Earth and also has a carbon dioxide atmosphere, which reduces the escape, but it also has a very low mass.

5.1 Hydrodynamic escape

An important loss mechanism that has been discussed extensively in the exoplanet literature is hydrodynamic escape. The term is usually used to refer to the outflow of atmospheres that are heated to such an extent that they flow away from the planets in the form of transonic Parker winds. This process does not take place in the solar system, mostly due to the Sun’s low activity. Hydrodynamic escape can be driven by the XUV spectrum of active stars (e.g. [186]), the energy held within the planet and atmosphere immediately after the circumstellar gas disk dissipates ([180]), or the photospheric emission of the star for very short period planets ([151]). Since atmospheres composed of lighter gases are lost more rapidly ([54]), these processes are most important for hydrogen dominated primordial atmospheres. For terrestrial planets with such atmospheres orbiting active stars, hydrodynamic escape is much stronger than stellar wind stripping ([106]). It is even possible that the pressure of the star’s wind suppresses the outflow and reduces the escape ([173];[194]). In hydrodynamical outflowing atmospheres, adiabatic cooling driven by the expansion of the gas has been shown to be an important thermal process ([186]; [141]).

A simple estimate for XUV-driven hydrodynamic escape rate can be achieved by considering the amount of energy in the star’s XUV field that is available to remove gas from the planet. If we assume a planet absorbs all XUV radiation that passes within a radius R_{XUV} of the planet’s center, the rate at which the planet absorbs stellar radiation is given by $\pi R_{\text{XUV}}^2 F_{\text{XUV}}$, where F_{XUV} is the XUV flux at the planet’s orbit. It is useful to define a parameter ϵ as the fraction of absorbed energy used to lift mass out of the gravitational potential well of the planet. This term is often called the ‘heating efficiency’ and this has caused it to

be confused with the fraction of absorbed XUV energy that is used to heat the gas. Much of the absorbed energy that is released as heat is irradiated back to space or ends up as kinetic energy of the escaping gas which is typically accelerated to speeds beyond that needed to leave the system. Since the energy required to remove a gravitationally bound mass m from near the surface of a planet is given by $GM_{\text{pl}}m/R_{\text{pl}}$, the rate at which a hydrodynamic flow removes energy is $(GM_{\text{pl}}/R_{\text{pl}})\dot{M}_{\text{at}}$, where $\dot{M}_{\text{at}}$ is the atmosphere's mass loss rate. This leads to an estimate of the mass loss rate given by

$$\dot{M}_{\text{at}} = \frac{\epsilon \pi R_{\text{XUV}}^2 R_{\text{pl}} F_{\text{XUV}}}{GM_{\text{pl}}}, \quad (63)$$

which is typically called the energy-limited escape equation ([201]).

This simple reasoning does not take into account several factors that can influence hydrodynamic losses. For instance, tidal effects are important for close-in planets ([53]). Since some of the energy is absorbed in the supersonic part of the wind and therefore cannot contribute to the loss rate and since the sonic point is closer to the planet for more rapidly outflowing atmospheres, the dependence of $\dot{M}_{\text{at}}$ on F_{XUV} should be weaker than linear (see Fig. 1 of [93]). Similarly, the loss of absorbed energy by radiative processes causes a weaker than linear $\dot{M}_{\text{at}}-F_{\text{XUV}}$ dependence for very highly irradiated planets ([141]; [148]). Also important for F_{XUV} is which parts of the XUV spectrum contribute to escape and this depends on the chemical composition of the atmosphere since the wavelength dependence for absorption are different for different chemical species. Since H and H₂ do not efficiently absorb radiation long ward of 112 nm, only the spectrum short-ward of this value is important. However, for Earth-like atmospheres, absorption by species such as O₂ and O₃ means that the relevant part of the spectrum extends beyond 200 nm, meaning more energy is available to drive hydrodynamic escape ([94]). For very strongly escaping hydrogen dominated atmospheres, it is possible that only the X-ray part of the spectrum is relevant if the EUV radiation is entirely absorbed above the sonic point ([148]). For a description of hydrodynamic escape that overcomes much of the limitations of the energy limited equation, see [111] and [112].

The thermal escape of an atmosphere and whether it is undergoing hydrodynamic escape depends primarily on the planet's mass, the mass of the atmospheric particles, and the temperature of the upper atmosphere and is often characterised by the Jeans escape parameter, λ_J . This is defined as the ratio of the potential energy to the thermal energy per particle at the exobase and is given by

$$\lambda_J = \frac{GM_{\text{pl}}m}{k_B T_{\text{exo}} R_{\text{exo}}}, \quad (64)$$

where m is the molecular mass of the gas, T_{exo} is the gas temperature, and R_{exo} is the exobase radius. This is also the ratio of the square of the escape velocity ($v_{\text{esc}}^2 = 2GM_{\text{pl}}/R_{\text{exo}}$) to the square of the most probable thermal velocity ($v_{\text{th}}^2 = 2k_B T_{\text{exo}}/m$) for particles at the exobase. This parameter specifies how gravitationally bound an atmosphere is to the planet and for $\lambda_J > 30$ the atmosphere can safely assumed to be hydrostatic, and for $\lambda_J < 15$, rapid thermal escape is expected to take place. For considering the escape of individual chemical species, m should be the mass of that species, and for considering if an atmosphere is undergoing hydrodynamic escape, m should be the average molecular mass of the gas. The name of this parameter comes from the fact that it is used in the expression for the Jeans escape rate, which for given species is given by

$$\dot{M}_J = 4\pi R_{\text{exo}}^2 n m v_{\text{th}} \frac{(1 + \lambda_J) e^{-\lambda_J}}{2\pi^{1/2}}, \quad (65)$$

where n is the number density at the exobase of the species of interest. For discussions on these processes, see [125] and [60].

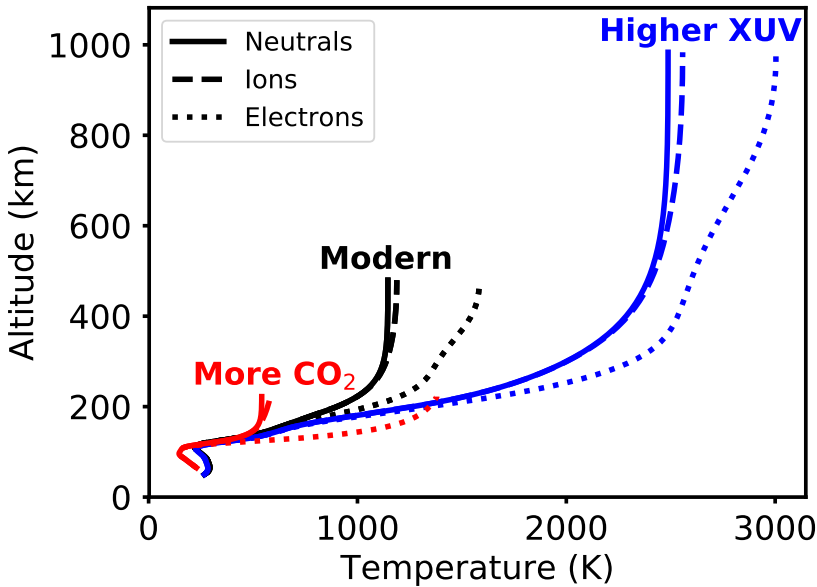


Figure 89: *Thermal structure of Earth’s mesosphere and thermosphere assuming modern conditions (black lines), the modern Earth’s atmosphere exposed to the higher XUV flux of the Sun approximately 2.5 Gyr ago (blue lines), and an Earth-like atmosphere with 100 times the present CO_2 composition exposed to the modern solar spectrum (red lines). The atmosphere models were calculated by [91].*

5.2 Upper atmospheres and active stars

While the modern Earth is a useful case to understand, we are more interested in this review in cases involving planets orbiting active stars since this is when the influences of stellar winds and atmospheric loss processes are most important. The higher XUV fluxes of more active stars causes the gas to be heated to much higher temperatures and to be ionised to a much greater degree. Several studies have explored what happens to the thermal and chemical structure of Earth’s thermosphere if it were present in its modern form at earlier times in the solar system’s history when the Sun was more active ([113]; [185]; [91], [94]). These studies have shown that the higher XUV flux of the younger more active Sun lead to thermospheres that are significantly hotter and more expanded, leading to more rapid escape at the upper boundary. An example of the effects of a higher XUV flux on the thermal structure of the upper atmosphere is shown in Fig. 89.

While we typically describe the atmospheres of the planets and moons in our solar system as hydrostatic, it is in fact always the case that some material is escaping from the upper atmosphere to space meaning that there is always a net upward flow of material throughout the atmosphere. For the Earth, this is typically much less than 1 cm s^{-1} at the exobase and is negligible. [185] studied the case of an input XUV spectrum with an EUV flux of twenty times the modern value and found that the thermosphere is heated to over 10^4 K and the exobase expands to an altitude of 10^5 km . This puts the exobase higher than the modern magnetopause of the Earth, meaning that the Earth’s thermosphere under such conditions would expand beyond the boundaries of the magnetosphere. Assuming only Jeans escape, they found that this causes upward flow speeds approaching 1 km s^{-1} which is high enough that adiabatic cooling of the expanding gas causes a negative temperature gradient in the

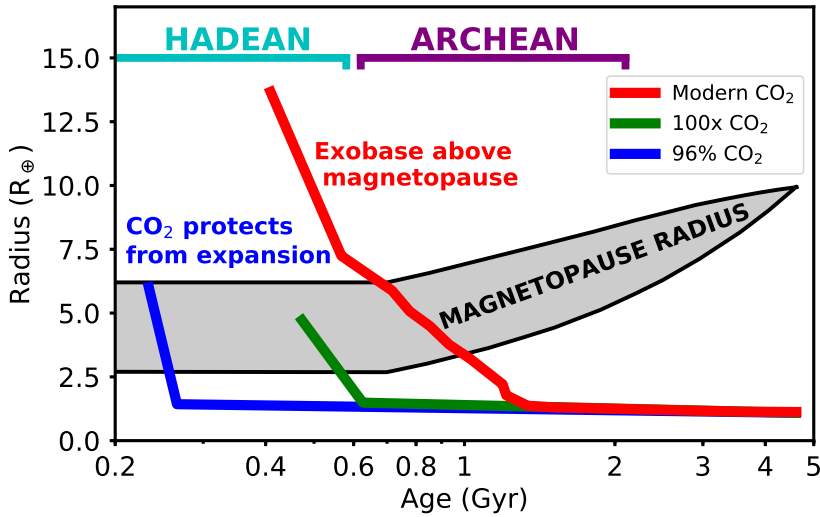


Figure 90: Estimates from [120] of the evolution of Earth's exobase altitude between 0.2 and 5 Gyr from the birth of the solar system assuming a modern atmospheric composition (red line), an atmosphere with 100x the modern CO_2 (green line), and an atmosphere that is 96% composed of CO_2 (blue line). The shaded area shows their range of estimates for the magnetopause radius. This figure shows that with a modern atmospheric composition, the Earth's thermosphere would have been directly exposed to the solar wind for the entire Hadean and the early Archean, and this can be avoided if a larger amount of CO_2 was present in the atmosphere.

upper thermosphere. Additional loss mechanisms, such as cold ion outflows and pick-up by stellar winds would cause additional adiabatic cooling in these cases.

[120] used the results of [185] to estimate the evolution of the Earth's thermospheric expansion and compared these with estimates the magnetosphere's size, as summarised in Fig. 90. The magnetosphere was likely smaller at earlier evolutionary phases due to stronger compression by the denser and faster early solar wind and a weaker intrinsic magnetic field ([184]). The total magnetospheric compression is uncertain and the range of scenarios calculated by [120] is shown as the shaded area in Fig. 90. They showed that with its current atmosphere, the Earth's thermosphere at ages younger than approximately 700 Myr would be expanded beyond the boundaries of the magnetosphere and be exposed to direct interactions with the solar wind. For their most active case, they estimated loss rates from ionisation and pick-up of atmospheric particles by the solar wind of 0.1 bar Myr^{-1} , meaning that the Earth's modern atmosphere would be removed in 10 Myr. Such rapid escape is unreasonable since it would likely imply that the Earth was without an atmosphere throughout the Hadean and early Archean, contrary to evidence for the presence during these times of an atmosphere, liquid surface water, and even life ([136]; [137]; [22]).

[94] considered an Earth with an Earth-like atmosphere exposed the much higher XUV flux of a very young rapidly rotating Sun and showed that the atmosphere is heated to such a high temperature that it flows away from the planet hydrodynamically. They found mass loss rates of approximately 10 bar Myr^{-1} , which is high enough to remove the entire modern atmosphere of the Earth in 0.1 Myr. They also showed that the gas is composed 20% of ions, which is several orders of magnitude more than the ionisation fraction in modern Earth's ionosphere.

5.3 The importance of atmospheric cooling

Consider again the results of [120] for the case of the Earth's atmosphere exposed to the higher XUV flux of the young Sun. It is interesting to speculate about what could have prevented the unrealistically high mass loss due to pick-up by the young solar wind that they found. An important process that likely protected Earth's atmosphere during the early period of high solar activity is upper atmosphere cooling by radiative emission to space. The most likely candidate for this cooling is infrared emission from CO_2 which is the most important coolant in Earth's upper atmosphere and is responsible for the formation of the mesosphere. Geochemical measurements show that the CO_2 abundance in Earth's atmosphere during and after the late Archean was orders of magnitude higher than it is in the modern atmosphere (e.g. [175]; [46]; [96]) and while such constraints on CO_2 for the mid Archean and earlier are not available, it is possible that CO_2 was the dominant component of the atmosphere during the Sun's highest activity phase.

The effects of having a CO_2 dominated atmosphere on the upper atmosphere of a planet can be seen well by comparing modern Earth and Venus. Venus' atmosphere is almost entirely CO_2 and despite receiving double the solar radiative flux as the Earth, its thermosphere is far colder and more contracted, with the exobase altitude being typically below 200 km and the exobase temperatures reaching only ~ 200 K ([77]; [67]). The effects of large changes in the CO_2 abundances in Earth's atmosphere were studied by [113] and [91] and an example of the effects of increasing the atmospheric CO_2 by a factor of 100 is shown in Fig. 89. The green and blue lines in Fig. 90 show the effects on the exobase radius of having 100x the modern CO_2 and a 96% CO_2 abundance respectively. The higher CO_2 abundances lead to much reduced exobase altitudes that are below the upper estimates for the magnetopause radius. The atmospheric loss rates from all important loss mechanisms in the cooler and less expanded cases are much lower.

The example of Earth's atmosphere evolution is useful for understanding the importance of atmospheric composition on the long term evolution of planetary atmospheres. While the influence of an intrinsic magnetic field on atmospheric losses is currently disputed and it is not clear if magnetospheres are indeed effective at protecting atmospheres from losses, it is clear that strong coolant molecules such as CO_2 are very effective at protecting atmospheres from interactions with stellar winds and from losses to space. Other important molecules for cooling include H_2O , NO , O , H , and H_3^+ . Cooling by H_2O is unimportant for modern Earth because the cold trap prevents significant amounts of water reaching the upper atmosphere. Under many conditions the cold trap does not form, such as those of the Earth prior to the great oxidation event 2.5 Gyr ago ([66]) and those of Earth-like planets orbiting low-mass stars ([65]), and this means that H_2O could play a much more important role. Cooling from species such as H_3^+ (see [135]) and H can be especially important for hydrogen dominated atmospheres such as those of gas giants ([141]; [174]; [23]). Note that not all cooling involves infrared emission: for example, hydrogen atoms cool by emitting in the Ly- α line which is part of the far ultraviolet.

5.4 The importance of the star's rotational evolution

Another important factor that can change the scenario for the Earth's thermosphere evolution shown in Fig. 90 is the rotation history of the Sun. [120] assumed a single time dependent decay law for the Sun's EUV emission based on the results of [165] to estimate the evolution of the thermospheric temperature. However, as we show in Fig. 78, the Sun's activity evolution depends on its initial rotation rate and different evolutionary tracks are possible. If the Sun was born as a slow rotator, its high-energy emission during the early phases of its

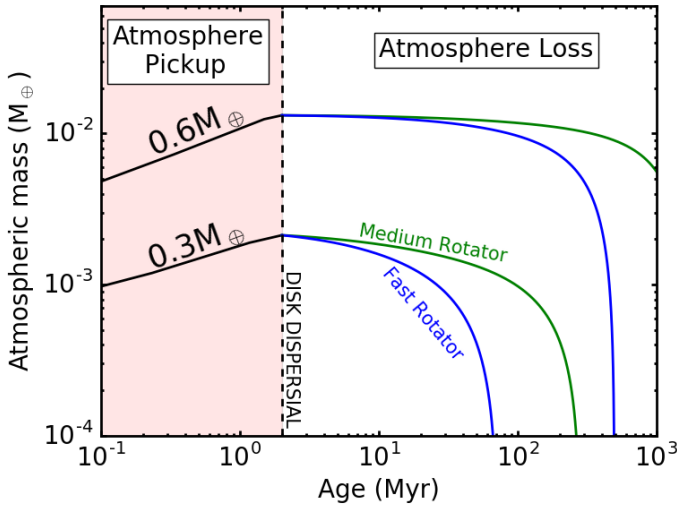


Figure 91: Example evolutionary tracks for the masses of hydrogen dominated primordial atmospheres on planets with masses of $0.3 M_{\oplus}$ and $0.6 M_{\oplus}$ orbiting solar mass stars at 1 AU. Cases for initially slowly and rapidly rotating stars are considered using the XUV evolution tracks of [189]. The pickup of the atmosphere during the disk phase is shown in the red shaded area and is based on the result of [181] and the loss after the disk has dissipated is calculated using the model developed by [93]. This shows the importance of a star's initial rotation rate for an atmosphere's evolution. While it is assumed in the calculations that the planets remain at the masses indicated for their entire lives, what matters most is the mass of the planet at the end of the gas disk phase.

evolution would have been lower than that assumed by [120]. It is likely also that the solar wind densities and speeds would have also been lower than assumed ([90]), though this is more speculative given our lack of understanding of the solar wind's evolution. Note that this section focuses on solar mass stars only and the differences between different rotation tracks is smaller for lower mass stars given their lower saturation thresholds for activity.

Given the link between stellar rotation and activity described in Section 2.3 and the diverse ways that rotation can evolve, the rotational evolution of the central star is an important factor that influences the long term evolution of atmospheres. This was explored for XUV driven hydrodynamic escape of primordial atmospheres in [93] and [110] who showed that planets orbiting initially rapidly rotating stars lose far more atmospheric gas than those orbiting initially slowly rotating stars, as demonstrated in Fig. 91. Depending on the initial rotation rate of the star, a planet can either lose an atmosphere entirely in the first billion years or keep it for its entire life, which has severe consequences for the conditions on the surfaces of these planets. The influence of different rotation tracks was studied for Earth-mass planets with water vapor atmospheres orbiting solar mass stars at 1 AU by [88] who showed that initially rapidly rotating stars erode far more atmospheric gas than slowly rotating stars. In the first billion years, a rapid rotator can remove several tens of Earth oceans of water (assuming a sufficiently large reservoir of water exists to be removed) whereas a slowly rotating star can only remove approximately 1 Earth ocean. Since water vapor is first photo-dissociated into atomic O and H (and ionized into O^+ and H^+) and the relative losses of H and O depend on

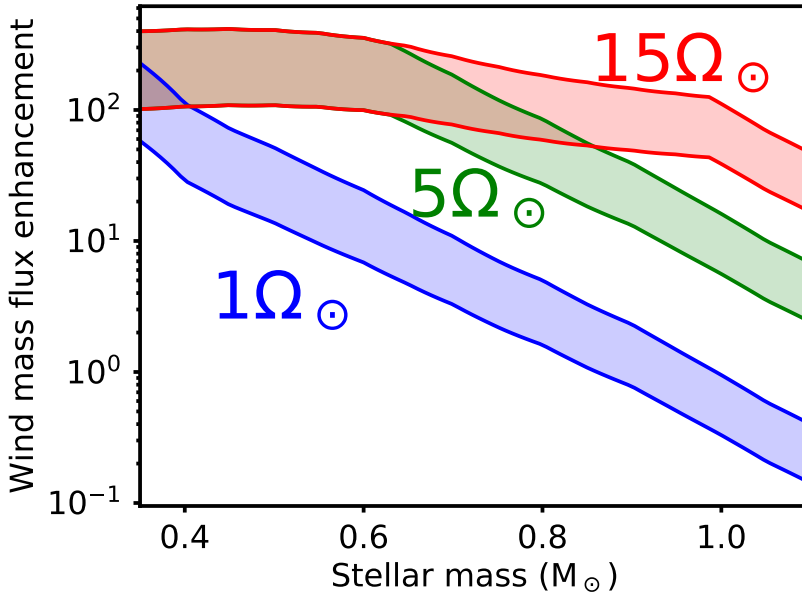


Figure 92: *Wind mass flux in the habitable zones of main-sequence stars as a function of stellar mass for three different rotation rates. The values are normalised to the mass flux of the solar wind at 1 AU and the ranges of values shown for each rotation rate show the fluxes between inner and outer habitable zone limits, defined as the moist and maximum greenhouse limits and calculated using the relations of [109].*

the loss rate, this has significant effects on how much O_2 remains on the atmosphere at the end of the escape phase and on the final atmosphere's composition.

The cases shown in Fig. 91 end at an age of 1 Gyr and we expect little additional loss at later ages since the activities of solar mass stars decay rapidly due to stellar wind driven spin-down. Even in the first billion years, the total amount of atmospheric escape is reduced by the decay in stellar rotation and activity as shown by [88] who considered also the effects of stars remaining at their maximum activity on losses from water vapor atmospheres. This is probably the most important effect that stellar winds have on the evolution of planetary atmospheres. By removing angular momentum from rapidly rotating young stars, winds prevent them from remaining highly active for their entire lives and therefore prevent the early period of extreme atmospheric escape from lasting until no volatile species are left to form an atmosphere. It is interesting to consider that despite driving a large number of important atmospheric loss processes, by spinning down their stars the net effect of stellar winds on the evolution of planetary atmospheres is likely one of atmospheric protection.

5.5 Stellar winds, CMEs, and habitable zone planets

Much of the motivation for studying how stellar winds influence the atmospheres of planets is our desire to understand questions related to planetary habitability and the likelihood that Earth-like surface conditions form on terrestrial planets outside our solar system. It is interesting for this purpose to consider the properties of stellar winds in the habitable zones (HZs) of stars with different masses and the effects on atmospheric losses. In Fig. 92, I show how

the mass flux in the wind in the HZ depends on stellar mass for three different rotation rates which I calculate using the constraints on wind mass loss derived from rotational evolution (Eqn. 61) as discussed in Section 3.1.1. Due to their much closer HZs, planets orbiting lower mass stars experience far higher mass fluxes when the stars are moderately or slowly rotating. Since lower mass stars experience longer periods of rapid rotation, we might expect that their HZ planets lose more atmospheric gas over longer time periods.

An important location in magnetised stellar winds is the Alfvén surface, which is the surface around the star at which the outflow speed of the wind becomes equal to the speed of Alfvén waves. The Alfvén surface is typically within a few tens of stellar radii and very non-spherical. [32] showed that for very low mass stars with HZs that are much closer to the stellar surface, the Alfvén surface can extend beyond the limits of the HZ, meaning that HZ planets experience both sub- and super-Alfvénic conditions over their orbits. This has an important effect on the magnetospheres of these planets since in the sub-Alfvénic regime, bow shocks do not form and the structure of the magnetosphere is very different. They also found that the extreme wind conditions experienced by these planets lead to very strong Joule heating of the atmospheres, which likely contributes to the expansion and loss of the atmosphere.

Interesting also is to consider the effects of CMEs which drive a large number of processes, and can even influence the lower atmospheres and climates of planets orbiting active stars. The effects of CMEs on HZ planets orbiting low mass stars was studied by [98] and [116] who argued that planets orbiting M dwarfs could be exposed to a continuous flux of CMEs which would compress their magnetospheres and drive a range of atmospheric loss processes. For such close HZs, CMEs often do not expand significantly before reaching the planet, which is also important for their influences on magnetospheres ([33]). Considering the case of the early Earth, [5] found that high-energy particles accelerated as a result of a large CME that propagate downwards through the atmosphere can significantly influence the chemical composition of the gas. The effect of a CME on loss processes in the exospheres of terrestrial planets was modelled by [106] who found stronger ionisation and pick-up of exospheric particles as CMEs pass the planet. Most importantly, these particles have sufficient energy to break the strong triple bonds of N_2 molecules, and the resulting dissociation products can form molecules such as N_2O , which is a strong greenhouse gas and might have influenced early Earth's climate.

6 Discussion

In this review, I tried to cover many of the aspects of stellar winds and planets that I consider important for the long term evolution of atmospheres. Discussions of this topic often concentrate on the interactions between winds and magnetospheres and the ability of planetary magnetic fields to protect atmospheres from losses. This often ignores many important processes especially in planetary thermospheres where stellar X-ray and ultraviolet radiation causes heating and expansion. Factors such as the XUV spectrum of the star and the chemical composition of the atmosphere are potentially more important for determining atmosphere–wind interactions than the properties of the wind and the strength of the planet's magnetic field. By cooling the upper atmosphere, chemical species such as carbon dioxide are very effective at protecting atmospheres from losses. It is also not clear what role planetary magnetic fields play in determining losses since they both weaken some loss mechanisms while strengthening others. While most loss mechanisms are partly driven by winds, the net effect of winds is probably one of atmospheric protection by spinning stars down and cause their activity levels to decay over evolutionary timescales. A full description of these processes studied in the context of stellar activity evolution is needed to understand the influences of stellar winds on planetary atmospheres.

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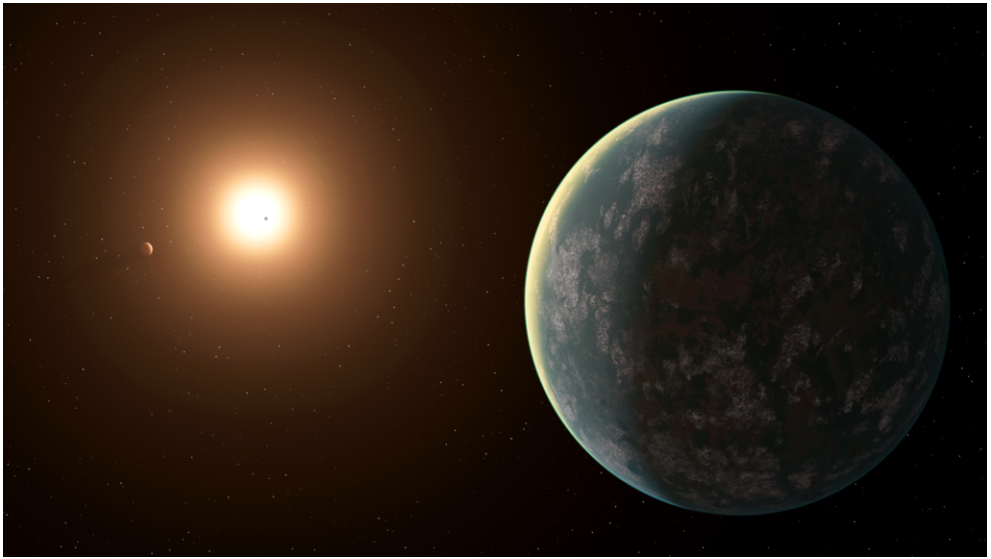
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Main ways in which stars influence the climate and surface habitability of their planets

by M. Turbet



Artist concept of the GJ357 system. GJ357b and c orbit in the potentially habitable zone of GJ357d. Copyright: Jack Madden. Licensed under the Creative Commons Attribution-Share Alike 4.0 International. Source https://commons.wikimedia.org/wiki/File:GJ_357_system.png

Main ways in which stars influence the climate and surface habitability of their planets

Martin Turbet^{1,*} and Franck Selsis²

¹Observatoire Astronomique de l'Université de Genève, Chemin Pegasi 51, 1290 Sauverny, Switzerland

²Laboratoire d'astrophysique de Bordeaux, Univ. Bordeaux, CNRS, B18N, Allée Geoffroy Saint-Hilaire, F-33615 Pessac, France

Abstract. We present a brief overview of the main effects by which a star will have an impact (positive or negative) on the surface habitability of planets in orbit around it. Specifically, we review how spectral, spatial and temporal variations in the incident flux on a planet can alter the atmosphere and climate of a planet and thus its habitability. For illustrative purposes, we emphasize the differences between planets orbiting solar-type stars and late M-stars. The latter are of particular interest as they constitute the first sample of potentially habitable exoplanets accessible for surface and atmospheric characterization in the coming years.

1 Introduction

From birth to death, the evolution of the nature of a planet (i.e. of its interior, surface and atmosphere) is intrinsically linked to that of its host star. Over time, as the nature of a planet evolves, its ability to offer all the necessary conditions for the emergence and development of life will thus evolve.

Based on our experience on Earth, we can identify some necessary (and certainly not sufficient) conditions for life to emerge and develop such as the availability of elements to make complex organic molecules (the famous "CHNOPS": carbon, hydrogen, nitrogen, oxygen, phosphorus and sulfur). Liquid water is considered an irreplaceable solvent for habitability (see [20] and references therein) but while it is known to exist in the interior of a large variety of planetary bodies (Europa, Ganymede, Callisto, Titan, Pluto, Triton, etc.; see [49] and references therein), planets that have liquid water on their surface have additional qualities. In fact, a source of low-entropy energy to initiate the irreversible chemical reactions necessary to initiate proto-metabolisms may require the exposition of liquid water to stellar radiations ([8, 57]). In the case of exoplanets, which will remain for a long time out of reach for in-situ exploration, identifying signs of life from interstellar distances (if possible) is likely to require photosynthetic processes and therefore surface liquid water. Lithoautotrophic life able to thrive without stellar light may indeed be unable to modify a whole planetary environment in an observable way ([31, 62]). In summary, the remote search for habitable planets outside the solar system boils down to the search for planets capable of maintaining liquid water on their surfaces ([30, 36]). This is why we focus in this manuscript on the main effects by which a star will have an impact on the ability of its planets to maintain surface liquid water.

*e-mail: martin.turbet@lmd.ipsl.fr

A first naive thing that comes to mind when thinking of the effects that a star may have on the habitability of its planets is the following thought experiment: what would happen to a planet without a star? Science fiction provides ideas of the situations in which this could happen. Although a planet around a black hole seems like an attractive configuration (see e.g., the *Interstellar* movie), black holes may lead to several warming mechanisms (e.g., by the blueshifting and beaming of incident radiation of the background stars on the planet ; see [64]) so that the system is not equivalent to a planet with no star. Better analogues are more likely rogue or free-floating planets, which are wandering planets without star. We now have direct evidence that such planets do exist by collecting the light directly emitted by the youngest, hottest and thus brightest of these objects ([7, 16, 18, 44, 48, 79]). Although these direct detections are limited so far to massive planets (several times the mass of Jupiter), we know these rogue planets exist in the terrestrial-mass regime too thanks to micro-lensing surveys ([55, 70]). The most likely formation scenario for these low-mass planets is that they have been ejected from their initial planetary system (see [38, 41, 50] and references therein). Rogue planets can maintain a subsurface ocean by the insulation of their internal heat flux by a thick layer of, for example, ice ([1]). They can even keep a surface temperature above 273 K under a dense enough H_2 atmosphere ([65, 68]) thanks to the opacity induced by H_2 - H_2 collisions ([10, 58, 60]). However, in the absence of stellar irradiation and given the challenge of sounding their atmosphere, the habitability of these objects is likely to remain an academic case for some time, unless there is a close encounter with the solar system ([1]). In short, having a star seems to be a necessary ingredient for a planet to be potentially habitable in a detectable way.

Therefore, we present below a brief overview of the dominant effects by which a star will have an impact on the habitability of planets in orbit around it. A more general list of all the effects that can affect the habitability of planets is provided in [66] and [53] but the main influence comes from the type of the host star, which strongly impacts both the possible lifespan of habitable conditions and our ability to remotely probe planetary environments. Stars more massive than $2.5 M_\odot$ live less than 1 Gigayear (Ga) and during this short life their luminosity increases dramatically offering conditions compatible for life on their planets for a fraction of this lifetime only. In addition, these massive stars are rare (less than 1 % of the galactic stellar population) and therefore distant while the detection and characterization of exoplanets around these bright stars is very challenging with our current observation techniques. At the other end of the stellar population, M stars represent ~ 75 % of the all stars, live tens to hundreds of Ga with a very stable luminosity (except for their earliest stage, which we will discuss below). Temperate terrestrial planets are frequent around these red-dwarfs and their characterization works far better with today's instruments. Therefore, along with Sun-like stars, they constitute a potential key population for habitable planets host-stars. To illustrate the ways in which a star may influence the habitability of its planets, we compare throughout the manuscript how the habitability of a planet varies – everything else being kept fixed – depending on whether it is orbiting a sun-like star or a late M-star. For the latter, we chose to use our closest neighbour, the star Proxima Centauri, as the archetype of a late M star ($M5.5V$, $T_{\text{eff}} = 3090K$) ; firstly because the star has been extremely well documented (see [61] and references therein) ; secondly because with the detection of a terrestrial-mass planet in temperate orbit ([2, 15, 34, 69]), it presents one of the most promising systems for our quest to find a habitable planet outside the solar system ([12, 43, 46, 73]).

The manuscript is organized as follows. Firstly, we review in Section 2 how the spatial and spectral distribution of the incident bolometric (mostly visible and near-infrared) flux affects the habitability of planets. Secondly, we discuss in Section 3 the influence of the far and mid-UV incident flux on the photochemistry, which can further affect the atmospheric

composition and thus the habitability of planets. Thirdly, we present in Section 4 how the X and EUV incident radiation can affect atmospheric escape, which can not only affect atmospheric composition but also endanger the very existence of an atmosphere around planets, which has severe consequences for their habitability. Fourthly, we show in Section 5 how the temporal evolution of the bolometric emission of a star (as well as its most energetic part, e.g. the XUV emission) impacts the habitability of planets around it.

2 Influence of the bolometric emission distribution on the climate

In this section, we briefly review how the climate and thus the habitability of a planet can be directly affected by the spectral type of its host star. The first naive effect is that the spectral type of the star changes its luminosity (for example, the luminosity of Proxima Cen is 0.15% that of the Sun) and thus the incident bolometric flux received by a planet at a fixed distance from the star. However, for a fixed amount of incident bolometric radiation received by the planet, the way in which stellar bolometric emission is distributed spectrally and spatially directly affects the way in which the surface, atmosphere and clouds absorb and reflect incident light, which changes the planet's surface temperature and thus the conditions of habitability.

The main factor in the spectral distribution of the incident bolometric flux on a planet is the effective temperature of the host star. According to Wien's law, the cooler the star, the more the bolometric emission is shifted to long wavelengths. As a matter of fact, while the peak of the Sun's emission ($T_{\text{eff}} = 5780\text{K}$) is $\sim 0.47 \mu\text{m}$, that of Proxima Centauri ($T_{\text{eff}} = 3090\text{K}$) is $\sim 1.1 \mu\text{m}$ (see Fig 93a). Note that due to the spectral shape of the Planck's law, the median of the stellar emission is shifted to higher wavelengths (compared to the peak). Median of solar emission is $\sim 0.73 \mu\text{m}$ while that of Proxima Centauri is $\sim 1.4 \mu\text{m}$ (see Fig 93b).

This has important consequences on a planet, because the surface, the atmosphere and the clouds absorb and/or scatter light differently at different wavelengths. Firstly, spectral variations of the surface albedo can have severe consequences on the stability of liquid water at a planetary surface. Ice and snow albedo are high at short wavelengths, and low at long wavelengths (see Fig 93c). As a result, ice and snow are much more reflective around a sun-like star than around a low-mass star, indicating that planets orbiting solar-type stars are more prone to glaciation (e.g., to be trapped in a snowball state) than planets orbiting low-mass stars ([29, 67, 74]). More generally, spectral variations in surface albedo, which have now been documented for many types of surface ([25]), can have a significant effect on a planet's climate ([51]). Secondly, Rayleigh scattering by the atmospheric gases can efficiently back-scatter (i.e. reflect back to space) incident stellar light, as illustrated by [32] for N_2 -dominated atmospheres, or by [36] for CO_2 -dominated atmospheres. Given that the Rayleigh scattering efficiency is $\propto \lambda^{-4}$ (see Fig 93d), the reflection of incoming stellar light by this process is much more efficient around a solar-type star than a low-mass star. Thirdly, direct absorption of incident stellar radiation by gases can vary greatly depending on the type of star. Most common gases lack strong absorption features in the optical wavelengths ([24]). This is illustrated in Fig 93e for H_2O absorption bands that absorb preferentially in the Proxima Cen emission wavelengths than those of the Sun. By combining the three ingredients mentioned above, we can conclude that – for a given incident flux – low-mass stars warm habitable planets more efficiently than sun-like stars do. This explains why the boundaries of the Habitable Zone are shifted towards lower incident stellar fluxes for low-mass stars (see e.g., Fig 8 in [36]).

Clouds may also have different properties whether they are exposed to a solar-like or a low-mass star incident radiation. This is illustrated in Figs. 94c, d and e which show the

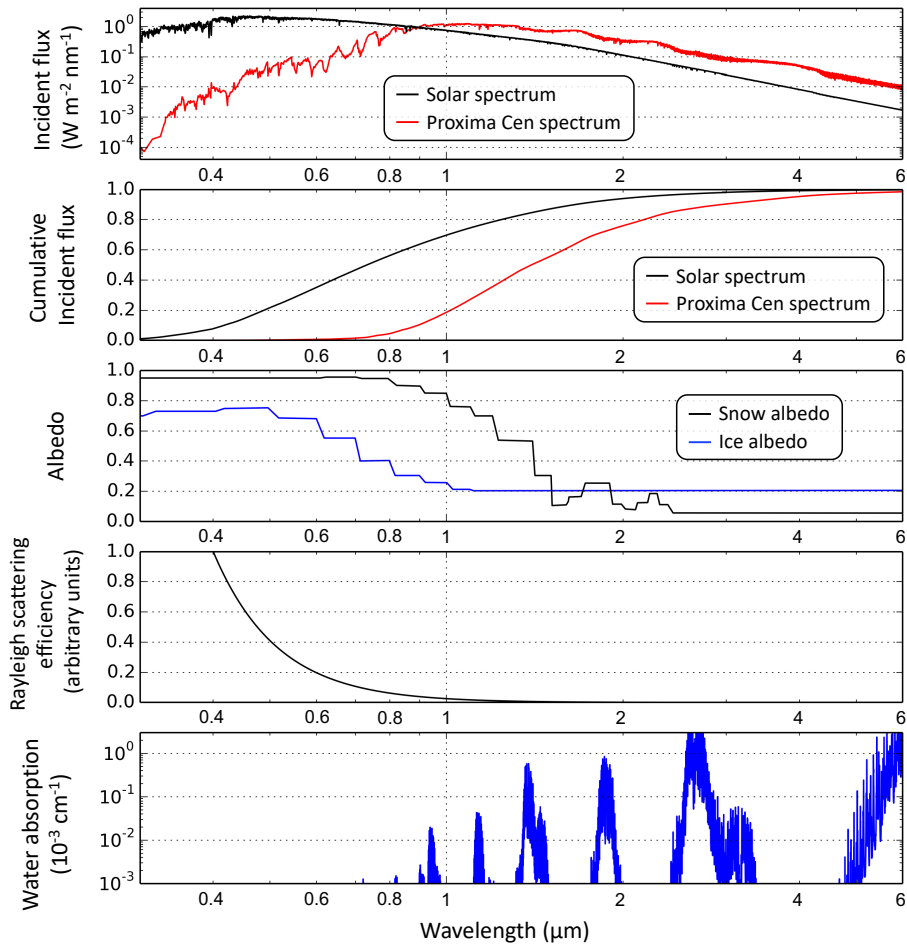


Figure 93: First panel: Incident flux spectra (at the top of the atmosphere) on a planet receiving a total of 1362 W m^{-2} (i.e. the solar constant on Earth ; note that the average flux received on the planet is four times lower, i.e. 340.5 W m^{-2}). The black curve is based on solar spectra reconstructed from [54] for $\lambda < 3\mu\text{m}$ and from [39] for $\lambda \geq 3\mu\text{m}$. The red curve is based on Proxima Centauri spectra reconstructed from [61]. Second panel: Cumulative incident flux spectra. Third panel: Albedo spectra for snow and water ice from [29]. Fourth panel: Rayleigh scattering cross-section spectrum (in arbitrary units). Fifth panel: Water vapor absorption spectrum, calculated at 200 K, 1 bar, and 0.1% of water vapor.

spectra of the key microphysical properties (extinction efficiency, single scattering albedo and asymmetry factor) of water ice clouds for several particle radii ([21]). Whether the presence of a particular type of ice particle or liquid water droplet serves to increase or decrease planetary albedo depends on the interplay among extinction efficiency, single scattering albedo and asymmetry factor ([21]). Clouds made of small particles (e.g., $0.1 \mu\text{m}$) are very reflective (asymmetry factor close to 0, single scattering albedo close to 1) for a wide range of stellar types. While clouds made of larger particles (e.g., $1 \mu\text{m}$ and larger) are still highly reflective,

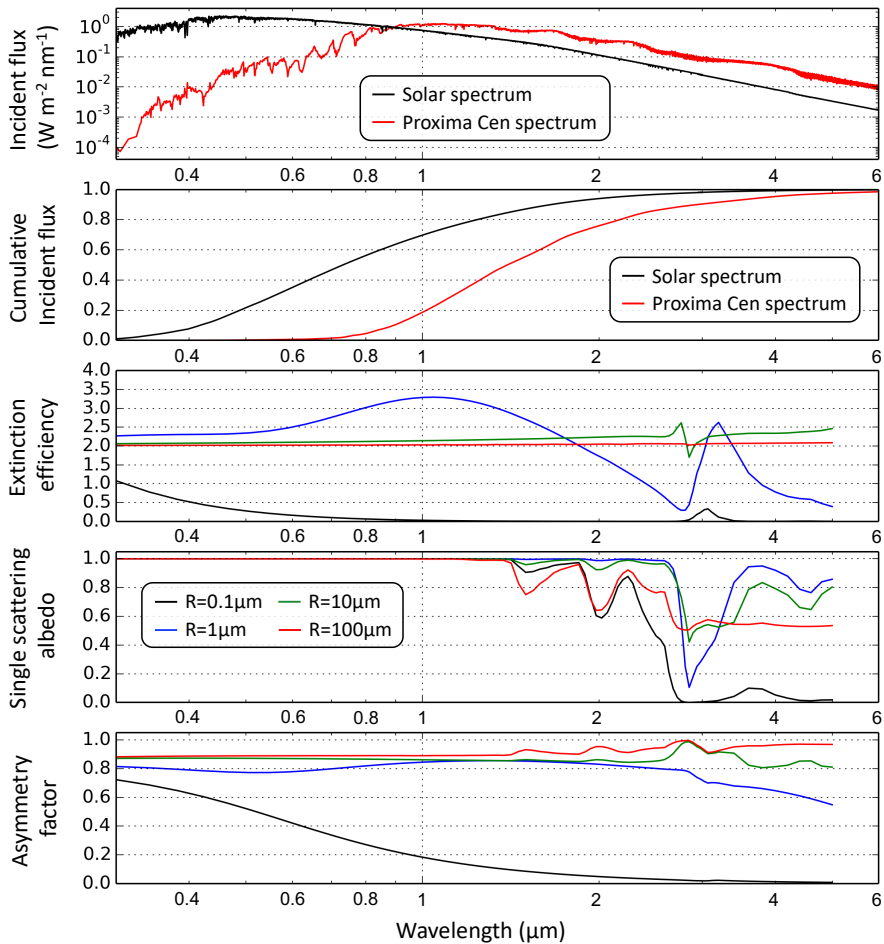


Figure 94: First panel: Incident flux spectra as in Fig. 93, panel 1. Second panel: Cumulative incident flux spectra as in Fig. 93, panel 2. Third panel: Water ice cloud extinction efficiencies for 4 spherical particle sizes (radius of 0.1, 1, 10 and $100\mu\text{m}$). Fourth panel: Water ice clouds single scattering albedos (i.e., the ratio of scattering efficiency to total extinction efficiency ; single-scattering albedo is equal to 1 if all is due to scattering, and equal to 0 if all particle extinction is due to absorption). Fifth panel: Water ice clouds asymmetry factors (i.e., the mean cosine of the scattering angle ; the asymmetry factor is 0 for isotropic radiation and ranges from -1 [negative values mean backward scattering is predominant] to 1 [positive values mean forward scattering is predominant].)

they can absorb an appreciable fraction of the incident flux near 1.5, 2 and $2.9\mu\text{m}$, which favors warming for planets around low-mass stars. However, overall, it has been shown that the differences between the different types of stars should be relatively small ([35]).

The main effect by which clouds may affect the climates of habitable planets differently depending on the type of host star is actually related to their spatial distribution. It has indeed been shown that planets orbiting in temperate orbits around low mass stars are prone to be

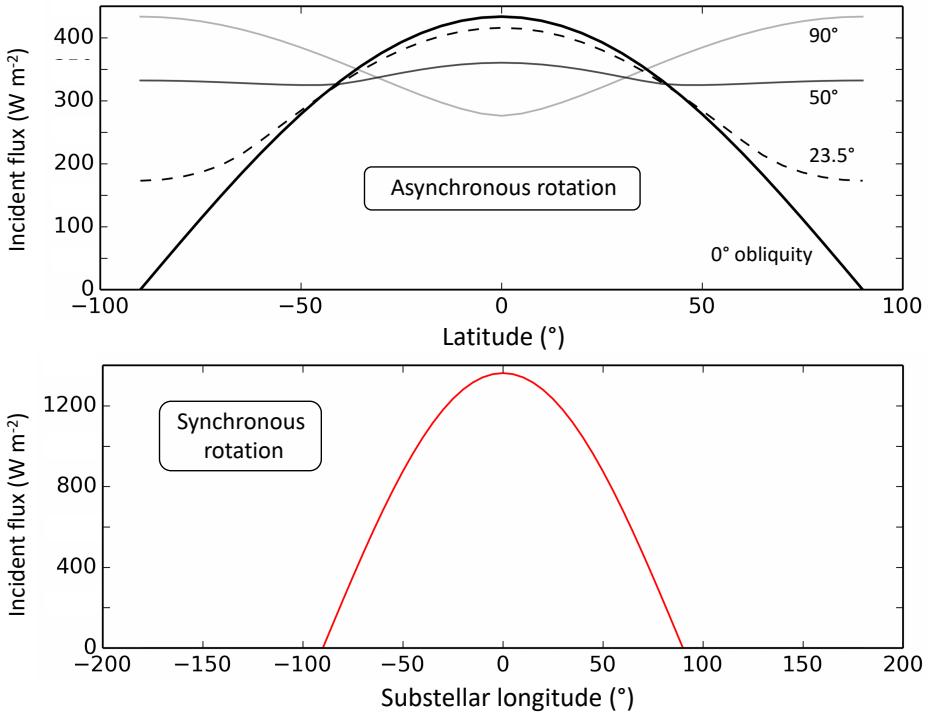


Figure 95: First panel: Zonal annual mean average of the incident flux on an asynchronously-rotating planet receiving a total of 1362 W m^{-2} (i.e. the solar constant on Earth), for four distinct obliquities (0° , 23.5° , 50° , 90°). Second panel: Incident flux on a synchronously-rotating planet (also receiving a total of 1362 W m^{-2}) versus substellar longitude. The zonal fluxes were calculated using equation 1 of [56].

locked in a state of synchronous rotation ([6, 17, 30]), for which one planet hemisphere is permanently facing its host star. This has severe consequences for the distribution of the insolation at the TOA (top of atmosphere) or surface of a planet, as illustrated in Fig. 95. While the solar flux is on average relatively evenly distributed on a fast-rotating planet (in particular at large obliquities ; see Fig. 95a), the stellar flux received on a synchronously-rotating planet is highly concentrated on one side of the planet (see Fig. 95b). It has been shown that on a planet covered with liquid water, this concentration of stellar flux leads to strong moist convection concentrated in the substellar region, which leads to the formation of a thick layer of reflecting clouds ([19, 37, 76, 77]). These clouds, because of their stabilizing effect (higher temperature means stronger convection, which means more clouds are formed, which means incoming stellar radiation is reflected more efficiently, which cools the planet), have a first-order effect on the climate and habitability of synchronously-rotating planets ([37, 76]), and thus on a significant fraction of planets orbiting low mass stars ([42]).

3 Influence of the UV emission on the photochemistry

The stellar Hydrogen Lyman- α (at 121.6 nm), far-ultraviolet ($\sim 122\text{-}200 \text{ nm}$) and mid-ultraviolet ($\sim 200\text{-}300 \text{ nm}$) spectral emissions are the main drivers of planetary photochem-

istry. UV ($\sim 100\text{-}300\text{ nm}$) photons can in fact photolyze (i.e. break) atmospheric molecules, forming radicals capable of generating multiple, complex chemical reactions. These photochemical reactions can deeply affect the atmospheric composition of a planet in a way that depends on initial composition and the total amount and spectral distribution of the UV stellar emission ([78]). Eventually, the photochemically-driven destruction or buildup of greenhouse gases in the atmosphere will affect planetary surface temperature and thus the surface habitability of the planet.

Each molecule has an UV absorption cross-section (see Fig. 96c) that, combined with the quantum yield (i.e. the probability that an absorbed photon actually breaks the molecule) and the UV incident spectral flux (Fig. 96a), gives the photolysis rate of the molecule. Depending on the spectral type of the host star, the photolysis rate of atmospheric molecules may – everything else being kept fixed – significantly vary. For illustration, the flux received by a planet (given a fixed bolometric incident flux) around a low-mass star is 2-3 orders of magnitude lower than that of a planet around a solar-type star in the main spectral region for ozone photodissociation (Hartley band ; near $\sim 250\text{ nm}$), a clue that oxygen and ozone chemistry must operate in a very different way around these two stars. More generally, the flux dichotomy near $\sim 150\text{ nm}$ (relative to its total bolometric luminosity, a M-star emit more photons below 150 nm , and much less photons above 150 nm) should have strong impact on the photochemistry of planetary atmospheres ([3, 52]).

Although in most atmospheres the impact of photochemistry is primarily to modify the relative abundances of minor species ([3, 52, 63]), there are some situations where photochemistry can entirely change the composition of an atmosphere ([23, 26, 45, 47]).

Finally, photochemistry can also lead to the formation of long carbonated chains suspended in the atmosphere, namely hazes. These radiatively active photochemical hazes can deeply affect the climate, either directly by absorbing and reflecting incoming stellar radiation in particular through Mie scattering, often resulting in surface cooling, or either indirectly through feedbacks on photochemistry e.g. by shielding molecules from incident UV radiation ([3]).

4 Influence of the X-EUV emission on atmospheric loss

X ($< 10\text{ nm}$) and Extreme-UV ($\sim 10\text{-}100\text{ nm}$) stellar emissions, also referred together as XUV, should play a fundamental role in atmospheric escape ([13, 40, 71]). XUV radiation heats the planetary thermosphere up to the exobase, which corresponds to the boundary where the atmosphere transitions from a collisional to a collisionless region. Hydrogen atoms are light enough to be lost to space when reaching the exobase even for a low XUV heating, like that of the Earth. The erosion of the water reservoir through this process is thus not controlled by the XUV flux but by the upward transport and photolysis of H_2O . Thanks to the cold trap of the tropopause that results in a H_2O -poor stratosphere, the erosion of Earth's water reservoir is negligible. When the XUV flux is sufficiently high, models predict that the heating powers an hydrodynamic expansion of the atmosphere, which is lost through a planetary wind ([28]). Even if hydrogen is the only species likely to be lost hydrodynamically with realistic XUV fluxes, other heavier species (e.g., oxygen) can also be accelerated through collisions and then be lost to space ([27]). In this hydrodynamical regime, the loss rate is controlled by the intensity of the XUV flux (as well as the availability of hydrogen atoms).

XUV incident flux on a planet can strongly vary depending on the type of host star, e.g. between 2-5 orders of magnitude depending on the exact wavelength between solar-type and low-mass stars (see Fig. 96a ; for a fixed bolometric incident flux), and depending also on the level of activity of the low-mass star. The total XUV flux of Proxima Cen is ~ 200 times higher than that of the Sun (see Fig. 96b), relatively to their total bolometric emission. This

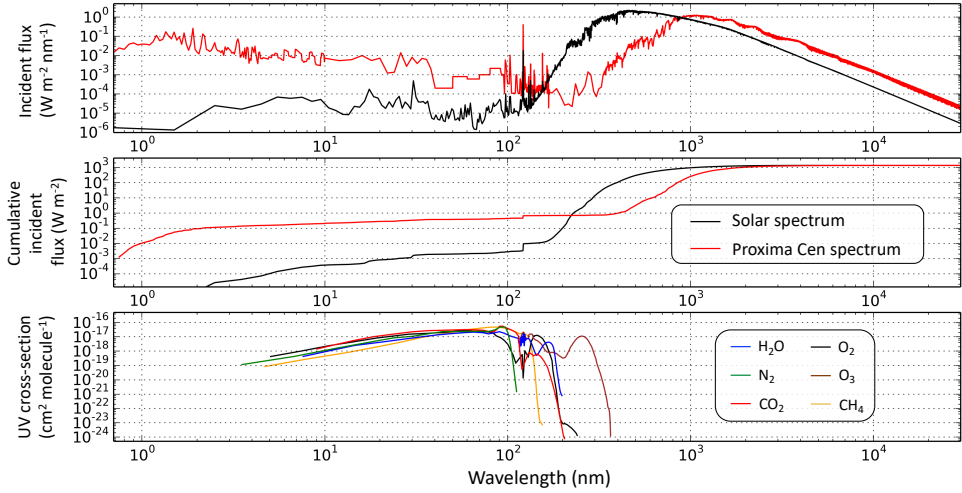


Figure 96: First panel: Incident flux spectra (at the top of the atmosphere) on a planet receiving a total of 1362 W m^{-2} (i.e. the solar constant on Earth). The black curve is based on solar spectra reconstructed from [54] for $\lambda < 3\mu\text{m}$ and from [39] for $\lambda \geq 3\mu\text{m}$. The red curve is based on Proxima Centauri spectra reconstructed from [61]. Second panel: Cumulative incident flux spectra. Third panel: UV cross-sections for common atmospheric species (H_2O , N_2 , CO_2 , O_2 , O_3 , CH_4) in terrestrial planetary atmospheres. The cross-sections were taken from [33].

means that the atmospheric escape rate can be significantly higher for planets around late M-stars than for Sun-like stars.

Most of stellar-driven atmospheric escape processes including XUV-driven escape (but also driven by the stellar wind drag or the photochemical escape) are also expected to be stronger around low-mass stars than solar-type stars (again, for a given total bolometric flux).

This is important information, given atmospheric escape can change the chemical composition or remove most of all of a planetary atmosphere. In the future, this threat against habitability around active late stars will have to be assessed with both observations/statistics of planetary atmospheres in various XUV/particle irradiation contexts and with robust modeling of the escape processes. The latter will require significant progress in order to treat in 3-D the transition between the fluid expanding atmosphere and the diluted particle regime where kinetic and electromagnetic interactions occurs far from LTE.

5 Influence of the temporal evolution of the stellar emissions

From birth to death, stars evolve in luminosity. They start their life with a super-luminous pre-main sequence phase during which they cool by radiating their initial accretion energy, then evolve on the main sequence during which their luminosity gradually increases as they fuse hydrogen into helium ([4, 5, 14]) and until they finally die. The duration of each of these two phases (pre-main sequence and main sequence), as well as the evolution of luminosity across them, has major consequences on the climate and habitability of the planets.

Firstly, the duration of the main sequence phase, which is all the longer the smaller the star, plays a key role in the stability of habitability conditions over time. While the duration of the main sequence of the Sun is a few billion years (but can be much shorter for higher mass stars), that of very low mass stars is expected to be much longer than the age of the Universe.

Secondly, while the pre-main sequence phase duration of solar-type stars is very short (~ 10 million years), that of very low mass stars can last for several hundreds of million years. During this phase, planets are exposed to a much larger incident flux, up to 10 x higher for planets around Proxima Cen (Fig. 97a). Planets that receive today the right incident flux to maintain surface liquid water were likely so hot (in the pre-main sequence phase of their host star) that they experienced a runaway greenhouse ([59]), with a thick steam atmosphere (assuming they formed with significant amounts of water) but no surface liquid water.

This is particularly critical for the habitability of planets around such stars, because water in the form of vapor is exposed to strong loss to space ([9, 11, 27, 47, 75]). Pre-main sequence phase can in fact potentially lead to complete water loss on a planet, depending on the initial reservoir. This is even more critical knowing that XUV heating, i.e. the main process by which the planet is likely to lose its atmosphere and its water, is also much more efficient during the pre-main sequence phase. This stems from the fact that the XUV radiation decreases even more rapidly than bolometric radiation during the pre-main-sequence phase (see Fig. 97b). Note that while the XUV flux emitted by a solar-like star may also have been very high in the past (at least for the fast rotator population ; see [72]), the cumulative XUV radiation (see Fig. 97c) is in general significantly larger around low-mass stars than sun-like stars. Combined with the long pre-main-sequence phase during which the planets are more irradiated than the runaway greenhouse threshold, and therefore during which water is exposed to atmospheric escape, this indicates that water loss is potentially much stronger on planets around late M-stars than around sun-like stars, which can have serious consequences for their habitability.

6 Conclusions

We briefly reviewed the main effects by which a star can have an impact on the habitability of planets in orbit around it. Firstly, spectral and spatial variations of the bolometric (i.e. visible and near-infrared) stellar emission can affect the way a planet's surface, atmosphere and clouds will absorb or reflect incident light. These changes can affect the climate of a planet (especially the surface temperature) which can thus impact its habitability. Secondly, variations in UV (far and mid-UV) incident fluxes on a planet can drive different photochemistry. This can potentially cause significant changes in the atmospheric composition, and thus the radiation balance of a planet, which can again strongly impact its habitability. Thirdly, cumulative variations in XUV (X and extreme-UV) incident fluxes on a planet can lead to various degrees of atmospheric erosion, potentially leading to dramatic effects (e.g. through the partial or complete loss of the atmosphere and the water) on the habitability.

Taking into account at the same time all the effects by which a star interacts with its planets in the same model, to simulate in a coherent way the evolution of the atmosphere and the habitability of the planets around any type of star, is one of the great challenges of the field. This ambitious scientific objective requires, among other things, a precise knowledge of the spectral properties (from infrared to XUV radiation) of the star at the time it is observed, but also of how these properties have evolved over time.

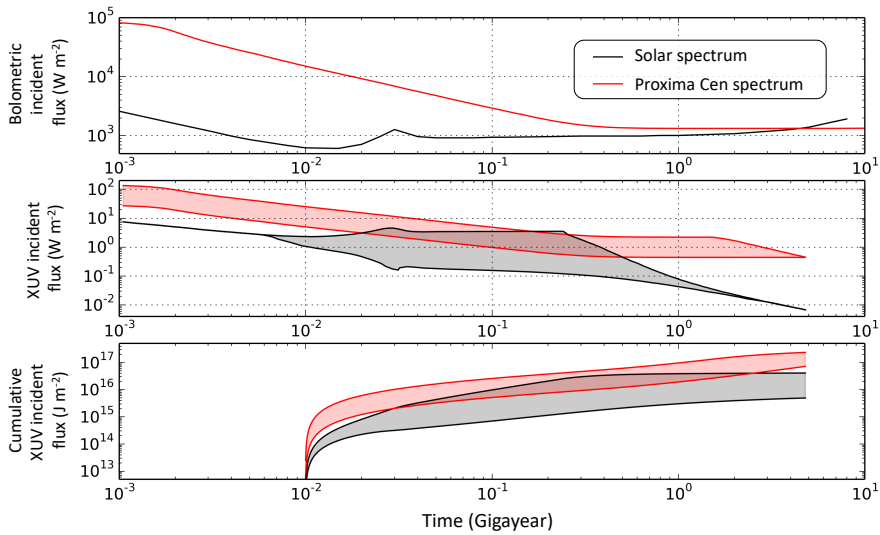


Figure 97: First panel: Temporal evolution of the incident flux (at the top of the atmosphere) on a planet receiving a total of 1362 W m^{-2} (i.e. the solar constant on Earth). The black curve is based on the solar evolution model of [5]. The red curve is based on the Proxima Centauri evolution model of [61]. Second panel: Temporal evolution of the XUV (10-120 nm) incident flux (at the top of the atmosphere) on a planet receiving a total of 1362 W m^{-2} (i.e. the solar constant on Earth). The black curves are based on the XUV emission solar models of [72] (based on [22]) for slow (lower curve) and fast (upper curve) rotator scenarios. The red curves are based on the two evolutionary scenarios of [61]: (1) Proxima Cen has spent its entire lifetime in the XUV emission saturation regime (lower curve); (2) Proxima Cen was in the saturation regime only for the first ~ 1.6 Gigayear of evolution (now aged of ~ 4.8 Gigayear). Third panel: Cumulative XUV incident flux starting 10 million years after the star formation, which is around the time the planets are expected to have formed.

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